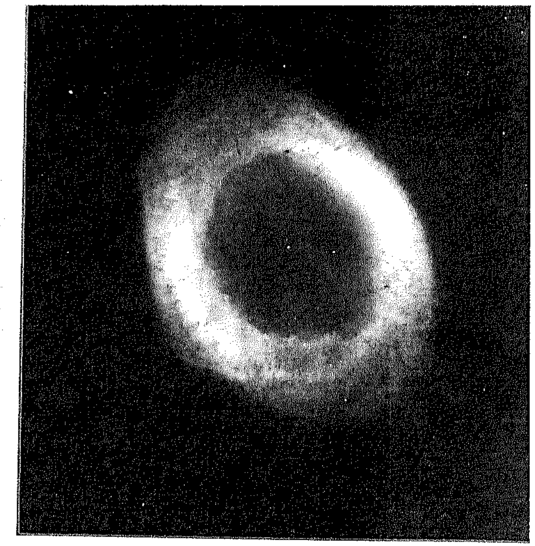
# 5 The Nature of Light

In the early 1800s, the French philosopher Auguste Comte argued that because the stars are so far away, humanity would never know their nature and composition. But the means to learn about the stars was already there for anyone to see—starlight. Just a few years after Comte's bold pronouncement, scientists began analyzing starlight to learn the very things that he had deemed unknowable.

We now know that atoms of each chemical element emit and absorb light at a unique set of wavelengths characteristic of that element alone. The red light in the accompanying image of a gas cloud in space is of a wavelength emitted by nitrogen and no other element; the particular green light in this image is unique to oxygen, and the particular blue light is unique to helium. The light from nearby planets, distant stars, and remote galaxies also has characteristic "fingerprints" that reveal the chemical composition of these celestial objects.

In this chapter we learn about the basic properties of light. Light has a dual nature: It has the properties of both waves and particles. The light emitted by an object depends upon the object's temperature; we can use this to determine the surface temperatures of stars. By studying the structure of atoms, we will learn why each element emits and absorbs light only at specific wavelengths and will see how astronomers determine what the atmospheres of planets and stars are made of. The motion of a light source also affects wavelengths, permitting us to deduce how fast stars and other objects are approaching or receding. These are but a few of the reasons why understanding light is a prerequisite to understanding the universe.



R I V U X G

The Ring Nebula is a shell of glowing gases surrounding a dying star. The spectrum of the emitted light reveals which gases are present. (Hubble Heritage Team, AURA/STScI/NASA)

## Learning Goals

*By reading the sections of this chapter, you will learn*

5-1 How we measure the speed of light

5-2 How we know that light is an electromagnetic wave

5-3 How an object's temperature is related to the radiation it emits

5-4 The relationship between an object's temperature and the amount of energy it emits

5-5 The evidence that light has both particle and wave aspects

5-6 How astronomers can detect an object's chemical composition by studying the light it emits

5-7 The quantum rules that govern the structure of an atom

5-8 The relationship between atomic structure and the light emitted by objects

5-9 How an object's motion affects the light we receive from that object

## 5-1 Light travels through empty space at a speed of 300,000 km/s

Galileo Galilei and Isaac Newton were among the first to ask basic questions about light. Does light travel instantaneously from one place to another, or does it move with a measurable speed? Whatever the nature of light, it does seem to travel swiftly from a source to our eyes. We see a distant event before we hear the accompanying sound. (For example, we see a flash of lightning before we hear the thunderclap.)

In the early 1600s, Galileo tried to measure the speed of light. He and an assistant stood at night on two hilltops a known distance apart, each holding a shuttered lantern. First, Galileo opened the shutter of

**The speed of light in a vacuum is a universal constant: It has the same value everywhere in the cosmos**

his lantern; as soon as his assistant saw the flash of light, he opened his own. Galileo used his pulse as a timer to try to measure the time between opening his lantern and seeing the light from his assistant's lantern. From the distance and time, he hoped to compute the speed at which the light had traveled to the distant hilltop and back.

Galileo found that the measured time failed to increase noticeably, no matter how distant the assistant was stationed. Galileo therefore concluded that the speed of light is too high to be measured by slow human reactions. Thus, he was unable to tell whether or not light travels instantaneously.

### The Speed of Light: Astronomical Measurements

The first evidence that light does *not* travel instantaneously was presented in 1676 by Olaus Rømer, a Danish astronomer. Rømer had been studying the orbits of the moons of Jupiter by carefully timing the moments when they passed into or out of Jupiter's shadow. To Rømer's surprise, the timing of these eclipses of Jupiter's moons seemed to depend on the relative positions of Jupiter and Earth. When Earth was far from Jupiter (that is, near conjunction; see Figure 4-6), the eclipses occurred several minutes later than when Earth was close to Jupiter (near opposition).

Rømer realized that this puzzling effect could be explained if light needs time to travel from Jupiter to Earth. When Earth is closest to Jupiter, the image of one of Jupiter's moons disappearing into Jupiter's shadow arrives at our telescopes a little sooner than it does when Jupiter and Earth are farther apart (Figure 5-1). The range of variation in the times at which such eclipses are observed is about 16.6 minutes, which Rømer interpreted as the length of time required for light to travel across the diameter of Earth's orbit (a distance of 2 AU). The size of Earth's orbit was not accurately known in Rømer's day, and he never actually calculated the speed of light. Today, using the modern value of 150 million kilometers for the astronomical unit, Rømer's method yields a value for the speed of light equal to roughly 300,000 km/s (186,000 mi/s).

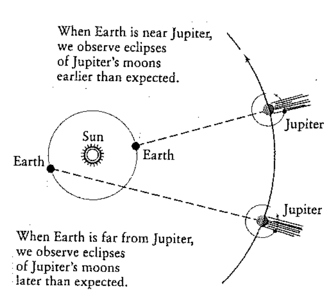



**Figure 5-1**

Rømer's Evidence That Light Does Not Travel Instantaneously The timing of eclipses of Jupiter's moons as seen from Earth depends on the Earth-Jupiter distance. Rømer correctly attributed this effect to variations in the time required for light to travel from Jupiter to Earth.

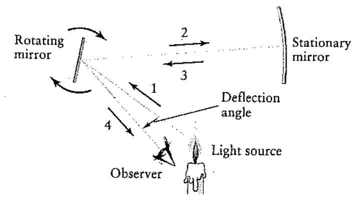



**Figure 5-2**

The Fizeau-Foucault Method of Measuring the Speed of Light Light from a light source (1) is reflected off a rotating mirror to a stationary mirror (2) and from there back to the rotating mirror (3). The ray that reaches the observer (4) is deflected away from the path of the initial beam because the rotating mirror has moved slightly while the light was making the round trip. The speed of light is calculated from the deflection angle and the dimensions of the apparatus.

### The Speed of Light: Measurements on Earth

Almost two centuries after Rømer, the speed of light was measured very precisely in an experiment carried out on Earth. In 1850, the French physicists Armand-Hippolyte Fizeau and Jean Foucault built the apparatus sketched in Figure 5-2. Light from a light source reflects from a rotating mirror toward a stationary mirror 20 meters away. The rotating mirror moves slightly while the light is making the round trip, so the returning light ray is deflected away from the source by a small angle. By measuring this angle and knowing the dimensions of their apparatus, Fizeau and Foucault could deduce the speed of light. Once again, the answer was very nearly 300,000 km/s.

The speed of light in a vacuum is usually designated by the letter *c* (from the Latin *celeritas,* meaning "speed"). The modern value is $c = 299{,}792.458$ km/s (186,282.397 mi/s). In most calculations you can use

$$c = 3.00 \times 10^5 \text{ km/s} = 3.00 \times 10^8 \text{ m/s}$$

The most convenient set of units to use for *c* is different in different situations. The value in kilometers per second (km/s) is often most useful when comparing *c* to the speeds of objects in space, while the value in meters per second (m/s) is preferred when doing calculations involving the wave nature of light (which we will discuss in Section 5-2).

**CAUTION!** Note that the quantity *c* is the speed of light *in a vacuum.* Light travels more slowly through air, water, glass, or any other transparent substance than it does in a vacuum. In our study of astronomy, however, we will almost always consider light traveling through the vacuum (or near-vacuum) of space.

The speed of light in empty space is one of the most important numbers in modern physical science. This value appears in many equations that describe atoms, gravity, electricity, and magnetism. According to Einstein's special theory of relativity, nothing can travel faster than the speed of light.



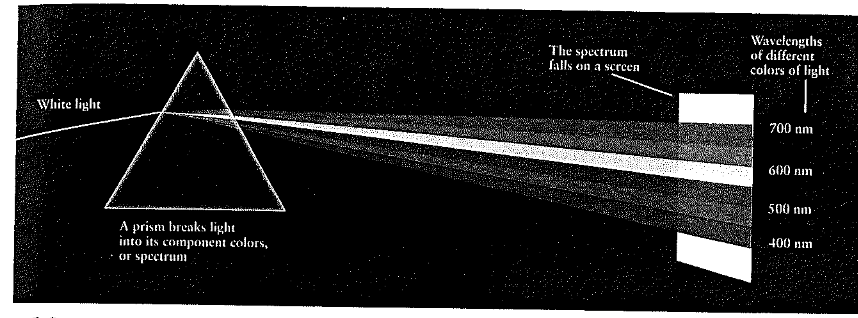



**Figure 5-3**

A Prism and a Spectrum When a beam of sunlight passes through a glass prism, the light is broken into a rainbow-colored band called a spectrum. The wavelengths of different colors of light are shown on the right (1 nm = 1 nanometer = $10^{-9}$ m).

## 5-2 Light is electromagnetic radiation and is characterized by its wavelength

Light is energy. This fact is apparent to anyone who has felt the warmth of the sunshine on a summer's day. But what exactly is light? How is it produced? What is it made of? How does it move through space? Scholars have struggled with these questions throughout history.

**Visible light, radio waves, and X rays are all the same type of wave: They differ only in their wavelength**

### Newton and the Nature of Color

The first major breakthrough in understanding light came from a simple experiment performed by Isaac Newton around 1670. Newton was familiar with what he called the "celebrated Phenomenon of Colours," in which a beam of sunlight passing through a glass prism spreads out into the colors of the rainbow (Figure 5-3). This rainbow is called a **spectrum** (plural **spectra**).

Until Newton's time, it was thought that a prism somehow added colors to white light. To test this idea, Newton placed a second prism so that just one color of the spectrum passed through it (Figure 5-4). According to the old theory, this should have caused a further change in the color of the light. But Newton found that each color of the spectrum was unchanged by the second prism; red remained red, blue remained

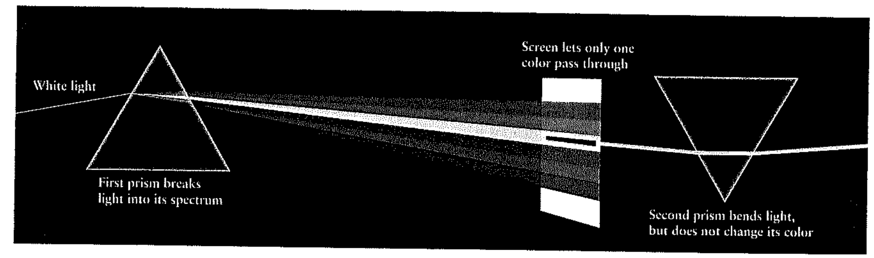



**Figure 5-4**

Newton's Experiment on the Nature of Light In a crucial experiment, Newton took sunlight that had passed through a prism and sent it through a second prism. Between the two prisms was a screen with a hole in it that allowed only one color of the spectrum to pass through. This same color emerged from the second prism. Newton's experiment proved that prisms do not add color to light but merely bend different colors through different angles. It also showed that white light, such as sunlight, is actually a combination of all the colors that appear in its spectrum.

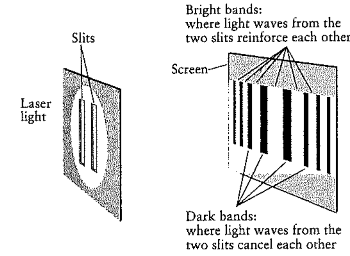
blue, and so on. He concluded that a prism merely separates colors and does not add color. Hence, the spectrum produced by the first prism shows that sunlight is a mixture of all the colors of the rainbow.

Newton suggested that light is composed of particles too small to detect individually. In 1678, however, the Dutch physicist and astronomer Christiaan Huygens proposed a rival explanation. He suggested that light travels in the form of waves rather than particles.

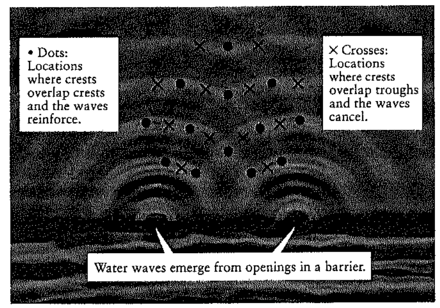
## Young and the Wave Nature of Light

Around 1801, Thomas Young in England carried out an experiment that convincingly demonstrated the wavelike aspect of light. He passed a beam of light through two thin, parallel slits in an opaque screen, as shown in Figure 5-5*a*. On a white surface some distance beyond the slits, the light formed a pattern of alternating bright and dark bands. Young reasoned that if a beam of light was a stream of particles (as Newton had suggested), the two beams of light from the slits should simply form bright images of the slits on the white surface. The pattern of bright and dark bands he observed is just what would be expected, however, if light had wavelike properties. An analogy with water waves demonstrates why.

ANALOGY Imagine ocean waves pounding against a reef or breakwater that has two openings (Figure 5-5*b*). A pattern of ripples is formed on the other side of the barrier as the waves come through the two openings and interfere with each other. At certain points, wave crests arrive simultaneously from the two openings. These reinforce each other and produce high waves. At other points, a crest from one opening meets a trough from the other opening. These cancel each other out, leaving areas of still water. This process of combining two waves also takes place in Young's double-slit experiment: The bright bands are regions where waves from the two slits reinforce each other, while the dark bands appear where waves from the two slits cancel each other.




(a) An experiment with light




(b) An analogous experiment with water waves

**Figure 5-5**

Young's Double-Slit Experiment **(a)** Thomas Young's classic double-slit experiment can easily be repeated in the modern laboratory by shining light from a laser onto two closely spaced parallel slits. Alternating dark and bright bands appear on a screen beyond the slits. **(b)** The intensity of light on the screen in (a) is analogous to the height of water waves that pass through a barrier with two openings. (The photograph shows this experiment with water waves in a small tank.) In certain locations, wave crests from both openings reinforce each other to produce extra high waves. At other locations a crest from one opening meets a trough from the other. The crest and trough cancel each other, producing still water. (Eric Schrempp/Photo Researchers)

## Maxwell and Light as an Electromagnetic Wave

The discovery of the wave nature of light posed some obvious questions. What exactly is "waving" in light? That is, what is it about light that goes up and down like water waves on the ocean? Because we can see light from the Sun, planets, and stars, light waves must be able to travel across empty space. Hence, whatever is "waving" cannot be any material substance. What, then, is it?

The answer came from a seemingly unlikely source—a comprehensive theory that described electricity and magnetism. Numerous experiments during the first half of the nineteenth century demonstrated an intimate connection between electric and magnetic forces. A central idea to emerge from these experiments is the concept of a *field*, an immaterial yet measurable disturbance of any region of space in which electric or magnetic forces are felt. Thus, an electric charge is surrounded by an electric field, and a magnet is surrounded by a magnetic field. Experiments in the early 1800s demonstrated that moving an electric charge produces a magnetic field; conversely, moving a magnet gives rise to an electric field.

In the 1860s, the Scottish mathematician and physicist James Clerk Maxwell succeeded in describing all the basic properties of electricity and magnetism in four equations. This mathematical achievement demonstrated that electric and magnetic forces are really two aspects of the same phenomenon, which we now call **electromagnetism.**

By combining his four equations, Maxwell showed that electric and magnetic fields should travel through space in the form of waves at a speed of $3.0 \times 10^5$ km/s—a value equal to the best available value for the speed of light. Maxwell's suggestion that



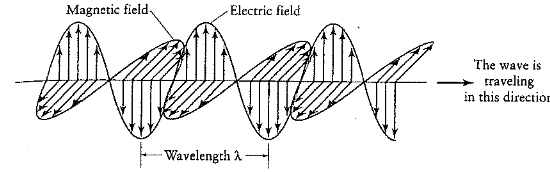



**Figure 5-6**

Electromagnetic Radiation All forms of light consist of oscillating electric and magnetic fields that move through space at a speed of $3.00 \times 10^5$ km/s = $3.00 \times 10^8$ m/s. This figure shows a "snapshot" of these fields at one instant. The distance between two successive crests, called the wavelength of the light, is usually designated by the Greek letter λ (lambda).

these waves do exist and are observed as light was soon confirmed by experiments. Because of its electric and magnetic properties, light is also called **electromagnetic radiation.**

CAUTION! You may associate the term *radiation* with radioactive materials like uranium, but this term refers to anything that radiates, or spreads away, from its source. For example, scientists sometimes refer to sound waves as "acoustic radiation." Radiation does not have to be related to radioactivity!

Electromagnetic radiation consists of oscillating electric and magnetic fields, as shown in Figure 5-6. The distance between two successive wave crests is called the **wavelength** of the light, usually designated by the Greek letter λ (lambda). No matter what the wavelength, electromagnetic radiation always travels at the same speed $c = 3.0 \times 10^5$ km/s = $3.0 \times 10^8$ m/s in a vacuum.

More than a century elapsed between Newton's experiments with a prism and the confirmation of the wave nature of light. One reason for this delay is that **visible light**, the light to which the human eye is sensitive, has extremely short wavelengths—less than a thousandth of a millimeter—that are not easily detectable. To express such tiny distances conveniently, scientists use a unit of length called the **nanometer** (abbreviated nm), where 1 nm = $10^{-9}$ m. Experiments demonstrated that visible light has wavelengths covering the range from about 400 nm for violet light to about 700 nm for red light. Intermediate colors of the rainbow like yellow (550 nm) have intermediate wavelengths, as shown in Figure 5-7. (Some astronomers prefer to measure wavelengths in *angstroms*. One angstrom, abbreviated Å, is one-tenth of a nanometer: 1 Å = 0.1 nm = $10^{-10}$ m. In these units, the wavelengths of visible light extend from about 4000 Å to about 7000 Å. We will not use the angstrom unit in this book, however.)

## Visible and Nonvisible Light

Maxwell's equations place no restrictions on the wavelength of electromagnetic radiation. Hence, electromagnetic waves could and should exist with wavelengths both longer and shorter than the 400–700 nm range of visible light. Consequently, researchers began to look for *invisible* forms of light. These are forms of electromagnetic radiation to which the cells of the human retina do not respond.

The first kind of invisible radiation to be discovered actually preceded Maxwell's work by more than a half century. Around 1800 the British astronomer William Herschel passed sunlight through a prism and held a thermometer just beyond the red end of the visible spectrum. The thermometer registered a temperature increase, indicating that it was being exposed to an invisible form of energy. This invisible energy, now called **infrared radiation,** was later realized to be electromagnetic radiation with wavelengths somewhat longer than those of visible light.

In experiments with electric sparks in 1888, the German physicist Heinrich Hertz succeeded in producing electromagnetic radiation with even longer wavelengths of a few centimeters or more. These are now known as **radio waves.** In 1895 another German physicist, Wilhelm Röntgen, invented a machine that produces electromagnetic radiation with wavelengths shorter than

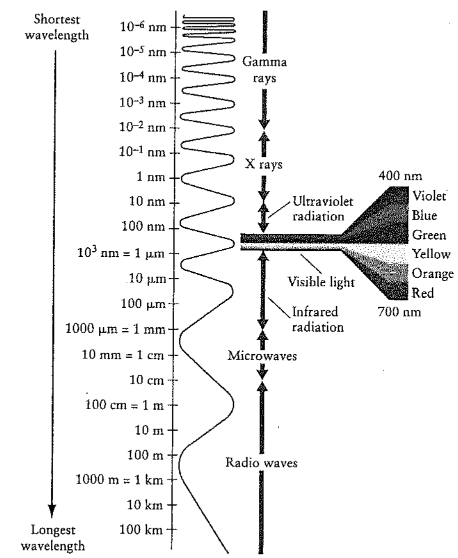



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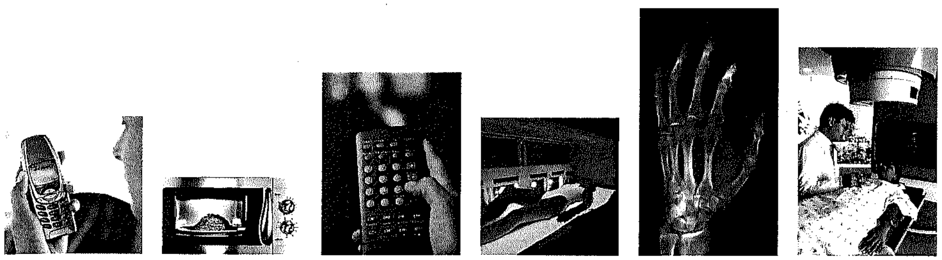
**Figure 5-7**

The Electromagnetic Spectrum The full array of all types of electromagnetic radiation is called the electromagnetic spectrum. It extends from the longest-wavelength radio waves to the shortest-wavelength gamma rays. Visible light occupies only a tiny portion of the full electromagnetic spectrum.



(a) Mobile phone: radio waves (b) Microwave oven: microwaves (c) TV remote: infrared light (d) Tanning booth: ultraviolet light (e) Medical imaging: X rays. (f) Cancer radiotherapy: gamma rays

Figure 5-8

Uses of Nonvisible Electromagnetic Radiation (a) A mobile phone is actually a radio transmitter and receiver. The wavelengths used are in the range 16 to 36 cm. (b) A microwave oven produces radiation with a wavelength near 10 cm. The water in food absorbs this radiation, thus heating the food. (c) A remote control sends commands to a television using a beam of infrared light. (d) Ultraviolet radiation in moderation gives you a suntan, but in excess can cause sunburn or skin cancer. (e) X rays can penetrate through soft tissue but not through bone, which makes them useful for medical imaging. (f) Gamma rays destroy cancer cells by breaking their DNA molecules, making them unable to multiply. (Ian Britton, Royalty-Free/Corbis, Michael Porsche/Corbis, Bill Lush/Taxi/Getty, Neil McAllister/Alamy, Edward Kinsman/Photo Researchers, Inc., Will and Deni McIntyre/Science Photo Library)

10 nm, now known as **X rays**. The X-ray machines in modern medical and dental offices are direct descendants of Röntgen's invention. Over the years radiation has been discovered with many other wavelengths.

Thus, visible light occupies only a tiny fraction of the full range of possible wavelengths, collectively called the **electromagnetic spectrum.** As Figure 5-7 shows, the electromagnetic spectrum stretches from the longest-wavelength radio waves to the shortest-wavelength gamma rays. Figure 5-8 shows some applications of nonvisible light in modern technology.

On the long-wavelength side of the visible spectrum, infrared radiation covers the range from about 700 nm to 1 mm. Astronomers interested in infrared radiation often express wavelength in *micrometers* or *microns,* abbreviated μm, where 1 μm = $10^{-3}$ mm = $10^{-6}$ m. **Microwaves** have wavelengths from roughly 1 mm to 10 cm, while radio waves have even longer wavelengths.

At wavelengths shorter than those of visible light, **ultraviolet radiation** extends from about 400 nm down to 10 nm. Next are X rays, which have wavelengths between about 10 and 0.01 nm, and beyond them at even shorter wavelengths are **gamma rays.** Note that the rough boundaries between different types of radiation are simply arbitrary divisions in the electromagnetic spectrum.

## Frequency and Wavelength

Astronomers who work with radio telescopes often prefer to speak of *frequency* rather than wavelength. The **frequency** of a wave is the number of wave crests that pass a given point in one second. Equivalently, it is the number of complete *cycles* of the wave that pass per second (a complete cycle is from one crest to the next). Frequency is usually denoted by the Greek letter $\nu$ (nu). The unit of frequency is the cycle per second, also called the *hertz* (abbreviated Hz) in honor of Heinrich Hertz, the physicist who first produced radio waves. For example, if 500 crests of a wave pass you in one second, the frequency of the wave is 500 cycles per second or 500 Hz.

In working with frequencies, it is often convenient to use the prefix *mega-* (meaning "million," or $10^6$, and abbreviated M) or *kilo-* (meaning "thousand," or $10^3$, and abbreviated k). For example, AM radio stations broadcast at frequencies between 535 and 1605 kHz (kilohertz), while FM radio stations broadcast at frequencies in the range from 88 to 108 MHz (megahertz).

The relationship between the frequency and wavelength of an electromagnetic wave is a simple one. Because light moves at a constant speed $c = 3 \times 10^8$ m/s, if the wavelength (distance from one crest to the next) is made shorter, the frequency must increase (more of those closely spaced crests pass you each second). The frequency $\nu$ of light is related to its wavelength $\lambda$ by the equation

**Frequency and wavelength of an electromagnetic wave**

$$\nu = \frac{c}{\lambda}$$

$\nu$ = frequency of an electromagnetic wave (in Hz)

$c$ = speed of light = $3 \times 10^8$ m/s

$\lambda$ = wavelength of the wave (in meters)

That is, the frequency of a wave equals the wave speed divided by the wavelength.

For example, hydrogen atoms in space emit radio waves with a wavelength of 21.12 cm. To calculate the frequency of this radiation, we must first express the wavelength in meters rather than centimeters: $\lambda = 0.2112$ m. Then we can use the above formula to find the frequency $\nu$:

$$\nu = \frac{c}{\lambda} = \frac{3 \times 10^8 \text{ m/s}}{0.2112 \text{ m}} = 1.42 \times 10^9 \text{ Hz} = 1420 \text{ MHz}$$

Visible light has a much shorter wavelength and higher frequency than radio waves. You can use the above formula to show that for yellow-orange light of wavelength 600 nm, the frequency is $5 \times 10^{14}$ Hz or 500 *million* megahertz!

While Young's experiment (Figure 5-5) showed convincingly that light has wavelike aspects, it was discovered in the early 1900s that light *also* has some of the characteristics of a stream of particles and waves. We will explore light's dual nature in Section 5-5.

## 5-3 An opaque object emits electromagnetic radiation according to its temperature

To learn about objects in the heavens, astronomers study the character of the electromagnetic radiation coming from those objects. Such studies can be very revealing because different kinds of electromagnetic radiation are typically produced in different ways. As an example, on Earth the most common way to generate radio waves is to make an electric current oscillate back and forth (as is done in the broadcast antenna of a radio station). By contrast, X rays for medical and dental purposes are usually produced by bombarding atoms in a piece of metal with fast-moving particles extracted from within other atoms. Our own Sun emits radio waves from near its glowing surface and X rays from its corona (see the photo that opens Chapter 3). Hence, these observations indicate the presence of electric currents near the Sun's surface and of fast-moving particles in the Sun's outermost regions. (We will discuss the Sun at length in Chapter 18.)

**As an object is heated, it glows more brightly and its peak color shifts to shorter wavelengths**

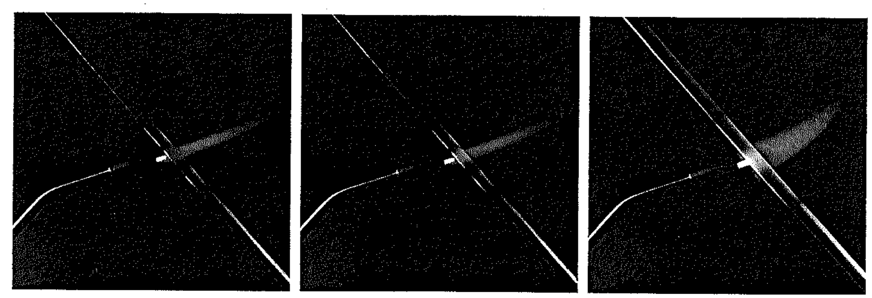
### Radiation from Heated Objects

TUTORIAL 5

The simplest and most common way to produce electromagnetic radiation, either on or off Earth, is to heat an object. The hot filament of wire inside an ordinary lightbulb emits white light, and a neon sign has a characteristic red glow because neon gas within the tube is heated by an electric current. In like fashion, almost all the visible light that we receive from space comes from hot objects like the Sun and the stars. The kind and amount of light emitted by a hot object tell us not only how hot it is but also about other properties of the object.

We can tell whether the hot object is made of relatively dense or relatively thin material. Consider the difference between a lightbulb and a neon sign. The dense, solid filament of a lightbulb makes white light, which is a mixture of all different visible wavelengths, while the thin, transparent neon gas produces light of a rather definite red color and, hence, a rather definite wavelength. For now we will concentrate our attention on the light produced by dense, opaque objects. (We will return to the light produced by gases in Section 5-6.) Even though the Sun and stars are gaseous, not solid, it turns out that they emit light with many of the same properties as light emitted by a hot, glowing, solid object.

Imagine a welder or blacksmith heating a bar of iron. As the bar becomes hot, it begins to glow deep red, as shown in Figure 5-9*a*. (You can see this same glow from the coils of a toaster or from an electric range turned on "high.") As the temperature rises further, the bar begins to give off a brighter orange light (Figure 5-9*b*). At still higher temperatures, it shines with a brilliant yellow light (Figure 5-9*c*). If the bar could be prevented from melt-



(a) Hot: glows deep red (b) Hotter: glows orange (c) Even hotter: glows yellow

Figure 5-9 R I V U X G

Heating a Bar of Iron This sequence of photographs shows how the appearance of a heated bar of iron changes with temperature. As the temperature increases, the bar glows more brightly because it radiates more energy. The color of the bar also changes because as the temperature goes up, the dominant wavelength of light emitted by the bar decreases. (©1984 Richard Megna/Fundamental Photographs)

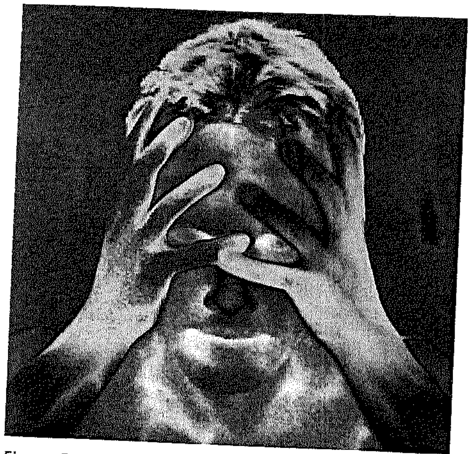


Figure 5-10 R I V U X G

An Infrared Portrait In this image made with a camera sensitive to infrared radiation, the different colors represent regions of different temperature. That this image is made from infrared radiation is indicated by the highlighted I in the wavelength tab. Red areas (like the man's face) are the warmest and emit the most infrared light, while blue-green areas (including the man's hands and hair) are at the lowest temperatures and emit the least radiation. (Dr. Arthur Tucker/Photo Researchers)

ing and vaporizing, at extremely high temperatures it would emit a dazzling blue-white light.

As this example shows, the amount of energy emitted by the hot, dense object and the dominant wavelength of the emitted radiation both depend on the temperature of the object. The hotter the object, the more energy it emits and the shorter the wavelength at which most of the energy is emitted. Colder objects emit relatively little energy, and this emission is primarily at long wavelengths.

These observations explain why you can't see in the dark. The temperatures of people, animals, and furniture are rather less than even that of the iron bar in Figure 5-9*a*. So, while these objects emit radiation even in a darkened room, most of this emission is at wavelengths greater than those of red light, in the infrared part of the spectrum (see Figure 5-7). Your eye is not sensitive to infrared, and you thus cannot see ordinary objects in a darkened room. But you can detect this radiation by using a camera that is sensitive to infrared light (Figure 5-10).

To better understand the relationship between the temperature of a dense object and the radiation it emits, it is helpful to know just what "temperature" means. The temperature of a substance is directly related to the average speed of the tiny **atoms**—the building blocks that come in distinct forms for each distinct chemical element—that make up the substance. (Typical atoms are about $10^{-10}$ m = 0.1 nm in diameter, or about 1/5000 as large as a typical wavelength of visible light.)

If something is hot, its atoms are moving at high speeds; if it is cold, its atoms are moving slowly. Scientists usually prefer to use the Kelvin temperature scale, on which temperature is measured in **kelvins** (K) upward from **absolute zero**. This is the coldest possible temperature, at which atoms move as slowly as possible (they can never quite stop completely). On the more familiar Celsius and Fahrenheit temperature scales, absolute zero (0 K) is −273°C and −460°F. Ordinary room temperature is 293 K, 20°C, or 68°F. Box 5-1 discusses the relationships among the Kelvin, Celsius, and Fahrenheit temperature scales.

Figure 5-11 depicts quantitatively how the radiation from a dense object depends on its Kelvin temperature. Each curve in this figure shows the intensity of light emitted at each wavelength by a dense object at a given temperature: 3000 K (the temperature at which molten gold boils), 6000K (the temperature of an iron-welding arc), and 12,000 K (a temperature found in special industrial furnaces). In other words, the curves show the spectrum of light emitted by such an object. At any temperature, a hot, dense object emits at all wavelengths, so its spectrum is a smooth, continuous curve with no gaps in it.

The shape of the spectrum depends on temperature, however. An object at relatively low temperature (say, 3000 K) has a low curve, indicating a low intensity of radiation. The **wavelength of maximum emission,** at which the curve has its peak and the emission of energy is strongest, is at a long wavelength. The higher the temperature, the higher the curve (indicating greater intensity) and the shorter the wavelength of maximum emission.

Figure 5-11 shows that for a dense object at a temperature of 3000 K, the wavelength of maximum emission is around 1000 nm (1 μm). Because this wavelength corresponds to the infrared range and is well outside the visible range, you might think that you cannot see the radiation from an object at this temperature. In fact, the

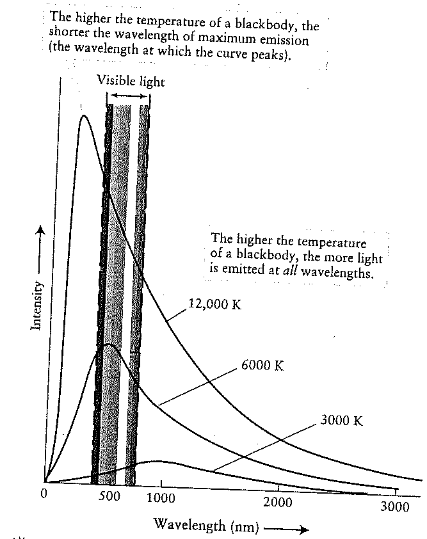



Figure 5-11

Blackbody Curves Each of these curves shows the intensity of light at every wavelength that is emitted by a blackbody (an idealized case of a dense object) at a particular temperature. The rainbow-colored band shows the range of visible wavelengths. The vertical scale has been compressed so that all three curves can be seen; the peak intensity for the 12,000-K curve is actually about 1000 times greater than the peak intensity for the 3000-K curve.



BOX 5-1 Tools of the Astronomer's Trade

## Temperatures and Temperature Scales

Three temperature scales are in common use. Throughout most of the world, temperatures are expressed in **degrees Celsius** (°C). The Celsius temperature scale is based on the behavior of water, which freezes at 0°C and boils at 100°C at sea level on Earth. This scale is named after the Swedish astronomer Anders Celsius, who proposed it in 1742.

Astronomers usually prefer the Kelvin temperature scale. This is named after the nineteenth-century British physicist Lord Kelvin, who made many important contributions to our understanding of heat and temperature. Absolute zero, the temperature at which atomic motion is at the absolute minimum, is −273°C in the Celsius scale but 0 K in the Kelvin scale. Atomic motion cannot be any less than the minimum, so nothing can be colder than 0 K; hence, there are no negative temperatures on the Kelvin scale. Note that we do *not* use degree (°) with the Kelvin temperature scale.

A temperature expressed in kelvins is always equal to the temperature in degrees Celsius plus 273. On the Kelvin scale, water freezes at 273 K and boils at 373 K. Water must be heated through a change of 100 K or 100°C to go from its freezing point to its boiling point. Thus, the "size" of a kelvin is the same as the "size" of a Celsius degree. When considering temperature changes, measurements in kelvins and Celsius degrees are the same. For extremely high temperatures the Kelvin and Celsius scales are essentially the same: for example, the Sun's core temperature is either $1.55 \times 10^7$ K or $1.55 \times 10^7$ °C.

The now-archaic Fahrenheit scale, which expresses temperature in degrees **Fahrenheit** (°F), is used only in the United States. When the German physicist Gabriel Fahrenheit introduced this scale in the early 1700s, he intended 100°F to represent the temperature of a healthy human body. On the Fahrenheit scale, water freezes at 32°F and boils at 212°F. There are 180 Fahrenheit degrees between the freezing and boiling points of water, so a degree Fahrenheit is only 100/180 = 5/9 as large as either a Celsius degree or a kelvin.

Two simple equations allow you to convert a temperature from the Celsius scale to the Fahrenheit scale and from Fahrenheit to Celsius:

$$T_F = \frac{9}{5}\,T_C + 32$$

$$T_C = \frac{5}{9}\,(T_F - 32)$$

$T_F$ = temperature in degrees Fahrenheit

$T_C$ = temperature in degrees Celsius

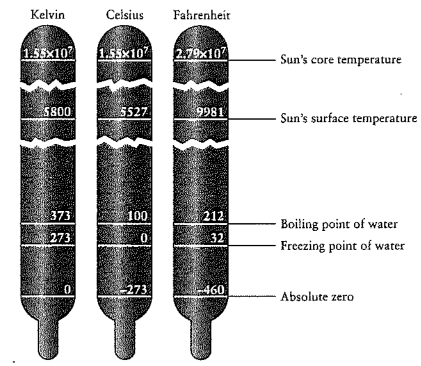
EXAMPLE: A typical room temperature is 68°F. We can convert this to the Celsius scale using the second equation:

$$T_C = \frac{5}{9}\,(68 - 32) = 20°C$$

To convert this to the Kelvin scale, we simply add 273 to the Celsius temperature. Thus,

$$68° = 20°C = 293\ K$$

The diagram displays the relationships among these three temperature scales.



glow from such an object *is* visible; the curve shows that this object emits plenty of light within the visible range, as well as at even shorter wavelengths.

The 3000-K curve is quite a bit higher at the red end of the visible spectrum than at the violet end, so a dense object at this temperature will appear red in color. Similarly, the 12,000-K curve has its wavelength of maximum emission in the ultraviolet part of the spectrum, at a wavelength shorter than visible light. But such a hot, dense object also emits copious amounts of visible light (much more than at 6000 K or 3000 K, for which the curves are lower) and thus will have a very visible glow. The curve for this temperature is higher for blue light than for red light, and so the color of a dense object at 12,000 K is a brilliant blue or blue-white. These conclusions agree with the color changes of a heated rod shown in Figure 5-9. The same principles apply to stars: A star that looks blue, such as Bellatrix in the constellation Orion (see Figure 2-2*a*), has a high surface temperature, while a red star such as Betelgeuse (Figure 2-2*a*) has a relatively cool surface.

These observations lead to a general rule:

***The higher an object's temperature, the more intensely the object emits electromagnetic radiation and the shorter the wavelength at which it emits most strongly.***

We will make frequent use of this general rule to analyze the temperatures of celestial objects such as planets and stars.

The curves in Figure 5-11 are drawn for an idealized type of dense object called a **blackbody.** A perfect blackbody does not reflect any light at all; instead, it absorbs all radiation falling on it. Because it reflects no electromagnetic radiation, the radiation that it does emit is entirely the result of its temperature. Ordinary objects, like tables, textbooks, and people, are not perfect blackbodies; they reflect light, which is why they are visible. A star such as the Sun, however, behaves very much like a perfect blackbody, because it absorbs almost completely any radiation falling on it from outside. The light emitted by a blackbody is called **blackbody radiation,** and the curves in Figure 5-11 are often called **blackbody curves.**

**CAUTION!** Despite its name, a blackbody does not necessarily *look* black. The Sun, for instance, does not look black because its temperature is high (around 5800 K), and so it glows brightly. But a room-temperature (around 300 K) blackbody would appear very black indeed. Even if it were as large as the Sun, it would emit only about 1/100,000 as much energy. (Its blackbody curve is far too low to graph in Figure 5-11.) Furthermore, most of this radiation would be at wavelengths that are too long for our eyes to perceive.

Figure 5-12 shows the blackbody curve for a temperature of 5800 K. It also shows the intensity curve for light from the Sun, as measured from above Earth's atmosphere. (This is necessary because Earth's atmosphere absorbs certain wavelengths.) The peak of both curves is at a wavelength of about 500 nm, near the middle of the visible spectrum. Note how closely the observed intensity curve for the Sun matches the blackbody curve. This is a strong indication that the temperature of the Sun's glowing surface is about 5800 K—a temperature that we can measure across a distance of 150 million kilometers! The close correlation between blackbody curves and the

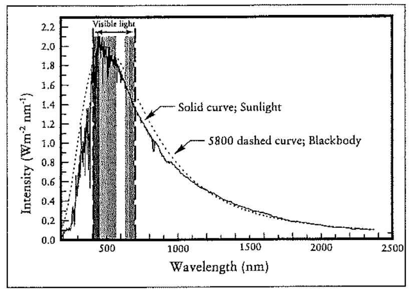



**Figure 5-12**

The Sun as a Blackbody This graph shows that the intensity of sunlight over a wide range of wavelengths (solid curve) is a remarkably close match to the intensity of radiation coming from a blackbody at a temperature of 5800 K (dashed curve). The measurements of the Sun's intensity were made above Earth's atmosphere (which absorbs and scatters certain wavelengths of sunlight). It's not surprising that the range of visible wavelengths includes the peak of the Sun's spectrum; the human eye evolved to take advantage of the most plentiful light available.

observed intensity curves for most stars is a key reason why astronomers are interested in the physics of blackbody radiation.

Blackbody radiation depends *only* on the temperature of the object emitting the radiation, not on the chemical composition of the object. The light emitted by molten gold at 2000 K is very nearly the same as that emitted by molten lead at 2000 K. Therefore, it might seem that analyzing the light from the Sun or from a star can tell astronomers the object's temperature but not what the star is made of. As Figure 5-12 shows, however, the intensity curve for the Sun (a typical star) is not precisely that of a blackbody. We will see later in this chapter that the *differences* between a star's spectrum and that of a blackbody allow us to determine the chemical composition of the star.

## 5-4 Wien's law and the Stefan-Boltzmann law are useful tools for analyzing glowing objects like stars

The mathematical formula that describes the blackbody curves in Figure 5-11 is a rather complicated one. But there are two simpler formulas for blackbody radiation that prove to be very useful in many branches of astronomy. They are used by astronomers who investigate the stars as well as by those who study the planets (which are dense, relatively cool objects that emit infrared radiation). One of these formulas relates the temperature of a blackbody to its wavelength of maximum emission, and the other relates the temperature to the amount of energy that the blackbody emits. These formulas, which we will use throughout this book, restate in precise mathematical terms the qualitative relationships that we described in Section 5-3.

**Two simple mathematical formulas describing blackbodies are essential tools for studying the universe**

### Wien's Law

Figure 5-11 shows that the higher the temperature (*T*) of a blackbody, the shorter its wavelength of maximum emission ($\lambda_{max}$). In 1893 the German physicist Wilhelm Wien used ideas about both heat and electromagnetism to make this relationship quantitative. The formula that he derived, which today is called **Wien's law,** is

**Wien's law for a blackbody**

$$\lambda_{max} = \frac{0.0029 \text{ K m}}{T}$$

$\lambda_{max}$ = wavelength of maximum emission of the object (in meters)

$T$ = temperature of the object (in kelvins)

According to Wien's law, the wavelength of maximum emission of a blackbody is inversely proportional to its temperature in kelvins. In other words, if the temperature of the blackbody doubles, its wavelength of maximum emission is halved, and vice versa. For example, Figure 5-11 shows blackbody curves for temperatures of 3000 K, 6000 K, and 12,000 K. From Wien's law, a blackbody with a temperature of 6000 K has a wavelength of maximum emission $\lambda_{max} = (0.0029 \text{ K m})/(6000 \text{ K}) = 4.8 \times 10^{-7}$ m = 480 nm, in the visible part of the electromagnetic spectrum. At 12,000 K, or twice the temperature, the blackbody has a wavelength of maximum emission half as great, or $\lambda_{max}$ = 240 nm, which is in the ultraviolet. At 3000 K, just half our original temperature, the value of $\lambda_{max}$ is twice the original value—960 nm, which is an infrared wavelength. You can see that these wavelengths agree with the peaks of the curves in Figure 5-11.

**CAUTION!** Remember that Wien's law involves the wavelength of maximum emission in *meters.* If you want to convert the wavelength to nanometers, you must multiply the wavelength in meters by ($10^9$ nm)/(1 m).

Wien's law is very useful for determining the surface temperatures of stars. It is not necessary to know how far away the star is, how large it is, or how much energy it radiates into space. All we need to know is the dominant wavelength of the star's electromagnetic radiation.

### The Stefan-Boltzmann Law

The other useful formula for the radiation from a blackbody involves the total amount of energy the blackbody radiates at all wavelengths. (By contrast, the curves in Figure 5-11 show how much energy a blackbody radiates at each individual wavelength.)

Energy is usually measured in **joules** (J), named after the nineteenth-century English physicist James Joule. A joule is the amount of energy contained in the motion of a 2-kilogram mass moving at a speed of 1 meter per second. The joule is a convenient unit of energy because it is closely related to the familiar **watt** (W): 1 watt is 1 joule per second, or 1 W = 1 J/s = 1 J $s^{-1}$. (The superscript −1 means you are dividing by that quantity.) For example, a 100-watt lightbulb uses energy at a rate of 100 joules per second, or 100 J/s. The energy content of food is also often measured in joules; in most of the world, diet soft drinks are labeled as "low joule" rather than "low calorie."

The amount of energy emitted by a blackbody depends both on its temperature and on its surface area. These characteristics make sense: A large burning log radiates much more heat than a burning match, even though the temperatures are the same. To consider the effects of temperature alone, it is convenient to look at the amount of energy emitted from each square meter of an object's surface in a second. This quantity is called the **energy flux** (*F*). Flux means "rate of flow," and thus *F* is a measure of how rapidly energy is flowing out of the object. It is measured in joules per square meter per second, usually written as J/m²/s or J $m^{-2}$ $s^{-1}$. Alternatively, because 1 watt equals 1 joule per second, we can express flux in watts per square meter (W/m², or W $m^{-2}$).

The nineteenth-century Irish physicist David Tyndall performed the first careful measurements of the amount of radiation emitted by a blackbody. (He studied the light from a heated platinum wire, which behaves approximately like a blackbody.) By analyzing Tyndall's results, the Slovenian physicist Josef Stefan deduced in 1879 that the flux from a blackbody is proportional to the fourth power of the object's temperature (measured in kelvins). Five years after Stefan announced his law, Austrian physicist Ludwig Boltzmann showed how it could be derived mathematically from basic assumptions about atoms and molecules. For this reason, Stefan's law is commonly known as the **Stefan-Boltzmann law.** Written as an equation, the Stefan-Boltzmann law is

**Stefan-Boltzmann law for a blackbody**

$$F = \sigma T^4$$

$F$ = energy flux, in joules per square meter of surface per second

$\sigma$ = a constant = $5.67 \times 10^{-8}$ W $m^{-2}$ $K^{-4}$

$T$ = object's temperature, in kelvins

The value of the constant $\sigma$ (the Greek letter sigma) is known from laboratory experiments.

The Stefan-Boltzmann law says that if you double the temperature of an object (for example, from 300 K to 600 K), then the energy emitted from the object's surface each second increases by a factor of $2^4 = 16$. If you increase the temperature by a factor of 10 (for example, from 300 K to 3000 K), the rate of energy emission increases by a factor of $10^4$ = 10,000. Thus, a chunk of iron at room temperature (around 300 K) emits very little electromagnetic radiation (and essentially no visible light), but an iron bar heated to 3000 K glows quite intensely.

Box 5-2 gives several examples of applying Wien's law and the Stefan-Boltzmann law to typical astronomical problems.

BOX 5-2 Tools of the Astronomer's Trade

## Using the Laws of Blackbody Radiation

The Sun and stars behave like nearly perfect blackbodies. Wien's law and the Stefan-Boltzmann law can therefore be used to relate the surface temperature of the Sun or a distant star to the energy flux and wavelength of maximum emission of its radiation. The following examples show how to do this.

EXAMPLE: The maximum intensity of sunlight is at a wavelength of roughly 500 nm = $5.0 \times 10^{-7}$ m. Use this information to determine the surface temperature of the Sun.

Situation: We are given the Sun's wavelength of maximum emission $\lambda_{max}$, and our goal is to find the Sun's surface temperature, denoted by $T_\odot$. (The symbol $\odot$ is the standard astronomical symbol for the Sun.)

Tools: We use Wien's law to relate the values of $\lambda_{max}$ and $T_\odot$.

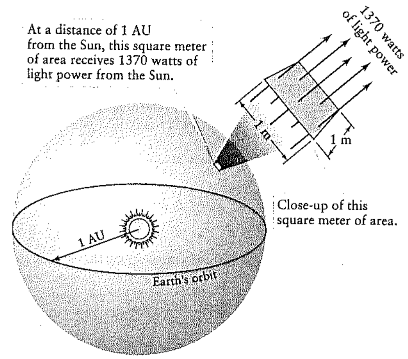
Answer: As written, Wien's law tells how to find $\lambda_{max}$ if we know the surface temperature. To find the surface temperature from $\lambda_{max}$, we first rearrange the formula, then substitute the value of $\lambda_{max}$:

$$T_\odot = \frac{0.0029 \text{ K m}}{\lambda_{max}} = \frac{0.0029 \text{ K m}}{5.0 \times 10^{-7} \text{ m}} = 5800 \text{ K}$$

Review: This temperature is very high by Earth standards, about the same as an iron-welding arc.

EXAMPLE: Using detectors above Earth's atmosphere, astronomers have measured the average flux of solar energy arriving at Earth. This value, called the **solar constant**, is equal to 1370 W $m^{-2}$. Use this information to calculate the Sun's surface temperature. (This calculation provides a check on our result from the preceding example.)

Situation: The solar constant is the flux of sunlight as measured at Earth. We want to use the value of the solar constant to calculate $T_\odot$.




Tools: It may seem that all we need is the Stefan-Boltzmann law, which relates flux to surface temperature. However, the quantity $F$ in this law refers to the flux measured at the Sun's surface, *not* at Earth. Hence, we will first need to calculate $F$ from the given information.

Answer: To determine the value of $F$, we first imagine a huge sphere of radius 1 AU with the Sun at its center, as shown in the figure. Each square meter of that sphere receives 1370 watts of power from the Sun, so the total energy radiated by the Sun per second is equal to the solar constant multiplied by the sphere's surface area. The result, called the **luminosity** of the Sun and denoted by the symbol $L_\odot$, is $L_\odot = 3.90 \times 10^{26}$ W. That is, in 1 second the Sun radiates $3.90 \times 10^{26}$ joules of energy into space. Because we know the size of the Sun, we can compute the energy flux (energy emitted per square meter per second) at its surface. The radius of the Sun is $R_\odot = 6.96 \times 10^8$ m, and the Sun's surface area is $4\pi R_\odot^2$.

Therefore, its energy flux $F_\odot$ is the Sun's luminosity (total energy emitted by the Sun per second) divided by the Sun's surface area (the number of square meters of surface):

$$F_\odot = \frac{L_\odot}{4\pi R_\odot^2} = \frac{3.90 \times 10^{26} \text{ W}}{4\pi (6.96 \times 10^8 \text{ m})^2} = 6.41 \times 10^7 \text{ W m}^{-2}$$

Once we have the Sun's energy flux $F_\odot$ we can use the Stefan-Boltzmann law to find the Sun's surface temperature $T_\odot$:

$$T_\odot^4 = F_\odot/\sigma = 1.13 \times 10^{15} \text{ K}^4$$

Taking the fourth root (the square root of the square root) of this value, we find the surface temperature of the Sun to be $T_\odot = 5800$ K.

Review: Our result for $T_\odot$ agrees with the value we computed in the previous example using Wien's law. Notice that the solar constant of 1370 W $m^{-2}$ is very much less than $F_\odot$, the flux at the Sun's surface. By the time the Sun's radiation reaches Earth, it is spread over a greatly increased area.

EXAMPLE: Sirius, the brightest star in the night sky, has a surface temperature of about 10,000 K. Find the wavelength at which Sirius emits most intensely.

Situation: Our goal is to calculate the wavelength of maximum emission of Sirius ($\lambda_{max}$) from its surface temperature $T$.

Tools: We use Wien's law to relate the values of $\lambda_{max}$ and $T$.

Answer: Using Wien's law,

$$\lambda_{max} = \frac{0.0029 \text{ K m}}{T} = \frac{0.0029 \text{ K m}}{10{,}000 \text{ K}}$$
$$= 2.9 \times 10^{-7} \text{ m} = 290 \text{ nm}$$

Review: Our result shows that Sirius emits light most intensely in the ultraviolet. In the visible part of the spectrum, it emits more blue light than red light (like the curve for 12,000 K in Figure 5-11), so Sirius has a distinct blue color.

EXAMPLE: How does the energy flux from Sirius compare to the Sun's energy flux?

Situation: To compare the energy fluxes from the two stars, we want to find the *ratio* of the flux from Sirius to the flux from the Sun.

Tools: We use the Stefan-Boltzmann law to find the flux from Sirius and from the Sun, which from the preceding examples have surface temperatures 10,000 K and 5800 K, respectively.

Answer: For the Sun, the Stefan-Boltzmann law is $F_\odot = \sigma T_\odot^4$, and for Sirius we can likewise write $F_* = \sigma T_*^4$, where the subscripts $\odot$ and * refer to the Sun and Sirius, respectively. If we divide one equation by the other to find the ratio of fluxes, the Stefan-Boltzmann constants cancel out and we get

$$\frac{F_*}{F_\odot} = \frac{T_*^4}{T_\odot^4} = \frac{(10{,}000 \text{ K})^4}{(5800 \text{ K})^4} = \left(\frac{10{,}000}{5800}\right)^4 = 8.8$$

Review: Because Sirius has such a high surface temperature, each square meter of its surface emits 8.8 times more energy per second than a square meter of the Sun's relatively cool surface. Sirius is actually a larger star than the Sun, so it has more square meters of surface area and, hence, its *total* energy output is *more* than 8.8 times that of the Sun.

## 5-5 Light has properties of both waves and particles

At the end of the nineteenth century, physicists mounted a valiant effort to explain all the characteristics of blackbody radiation. To this end they constructed theories based on Maxwell's description of light as electromagnetic waves. But all such theories failed to explain the characteristic shapes of blackbody curves shown in Figure 5-11.

### The Photon Hypothesis

In 1900, however, the German physicist Max Planck discovered that he could derive a formula that correctly described blackbody curves if he made certain assumptions. In 1905 the great German-born physicist Albert Einstein realized that these assumptions implied a radical new view of the nature of light. One tenet of this new view is that electromagnetic energy is emitted in discrete, particle-like packets, or light *quanta* (the plural of *quantum*, from a Latin word meaning "how much"). The second tenet is that the energy of each light quantum—today called a **photon**—is related to the wavelength of light: the greater the wavelength, the lower the energy of a photon associated with that wavelength. Thus, a photon of red light (wavelength $\lambda$ = 700 nm) has less energy than a photon of violet light ($\lambda$ = 400 nm). In this picture, light has a dual personality; it behaves as a stream of particlelike photons, but each photon has wavelike properties. In this sense, the best answer to the question "Is light a wave or a stream of particles?" is "Yes!"

**The revolutionary photon concept is necessary to explain blackbody radiation and the photoelectric effect**

It was soon realized that the photon hypothesis explains more than just the detailed shape of blackbody curves. For example, it explains why only ultraviolet light causes suntans and sunburns. The reason is that tanning or burning involves a chemical reaction in the skin. High-energy, short-wavelength ultraviolet photons can trigger these reactions, but the lower-energy, longer-wavelength photons of visible light cannot. Similarly, normal photographic film is sensitive to visible light but not to infrared light; a long-wavelength infrared photon does not have enough energy to cause the chemical change that occurs when film is exposed to the higher-energy photons of visible light.

Another phenomenon explained by the photon hypothesis is the **photoelectric effect**. In this effect, a metal plate is illuminated by a light beam. If ultraviolet light is used, tiny negatively charged particles called **electrons** are emitted from the metal plate. (We will see in Section 5-7 that the electron is one of the basic particles of the atom.) But if visible light is used, no matter how bright, no electrons are emitted.

Einstein explained this behavior by noting that a certain minimum amount of energy is required to remove an electron from the metal plate. The energy of a short-wavelength ultraviolet photon is greater than this minimum value, so an electron that absorbs a photon of ultraviolet light will have enough energy to escape from the plate. But an electron that absorbs a photon of visible light, with its longer wavelength and lower energy, does not gain enough energy to escape and so remains within the metal. Einstein and Planck both won Nobel prizes for their contributions to understanding the nature of light.

## The Energy of a Photon

The relationship between the energy $E$ of a single photon and the wavelength of the electromagnetic radiation can be expressed in a simple equation:

**Energy of a photon (in terms of wavelength)**

$$E = \frac{hc}{\lambda}$$

$E$ = energy of a photon

$h$ = Planck's constant

$c$ = speed of light

$\lambda$ = wavelength of light

The value of the constant $h$ in this equation, now called *Planck's constant,* has been shown in laboratory experiments to be

$$h = 6.625 \times 10^{-34} \text{ J s}$$

The units of $h$ are joules multiplied by seconds, called "joule-seconds" and abbreviated J s.

Because the value of $h$ is so tiny, a single photon carries a very small amount of energy. For example, a photon of red light with wavelength 633 nm has an energy of only $3.14 \times 10^{-19}$ J (Box 5-3), which is why we ordinarily do not notice that light comes in the form of photons. Even a dim light source emits so many photons per second that it seems to be radiating a continuous stream of energy.

The energies of photons are sometimes expressed in terms of a small unit of energy called the **electron volt** (eV). One electron volt is equal to $1.602 \times 10^{-19}$ J, so a 633-nm photon has an energy of 1.96 eV. If energy is expressed in electron volts, Planck's constant is best expressed in electron volts multiplied by seconds, abbreviated eV s:

$$h = 4.135 \times 10^{-15} \text{ eV s}$$

Because the frequency $\nu$ of light is related to the wavelength $\lambda$ by $\nu = c/\lambda$, we can rewrite the equation for the energy of a photon as

**Energy of a photon (in terms of frequency)**

$$E = h\nu$$

$E$ = energy of a photon

$h$ = Planck's constant

$\nu$ = frequency of light

The equations $E = hc/\lambda$ and $E = h\nu$ are together called **Planck's law.** Both equations express a relationship between a particlelike property of light (the energy $E$ of a photon) and a wavelike property (the wavelength $\lambda$ or frequency $\nu$).

The photon picture of light is essential for understanding the detailed shapes of blackbody curves. As we will see, it also helps to explain how and why the spectra of the Sun and stars differ from those of perfect blackbodies.

**BOX 5-3** Astronomy Down to Earth

### Photons at the Supermarket

A beam of light can be regarded as a stream of tiny packets of energy called photons. The Planck relationships $E = hc/\lambda$ and $E = h\nu$ can be used to relate the energy $E$ carried by a photon to its wavelength $\lambda$ and frequency $\nu$.

As an example, the laser bar-code scanners used at stores and supermarkets emit orange-red light of wavelength 633 nm. To calculate the energy of a single photon of this light, we must first express the wavelength in meters. A nanometer (nm) is equal to $10^{-9}$ m, so the wavelength is

$$\lambda = (633 \text{ nm})\left(\frac{10^{-9}\text{ m}}{1 \text{ nm}}\right) = 633 \times 10^{-9} \text{ m} = 6.33 \times 10^{-7} \text{ m}$$

Then, using the Planck formula $E = hc/\lambda$, we find that the energy of a single photon is

$$E = \frac{hc}{\lambda} = \frac{(6.625 \times 10^{-34} \text{ J s})(3 \times 10^{8} \text{ m/s})}{6.33 \times 10^{-7} \text{ m}} = 3.14 \times 10^{-19} \text{ J}$$

This amount of energy is very small. The laser in a typical bar-code scanner emits $10^{-3}$ joule of light energy per second, so the number of photons emitted per second is

$$\frac{10^{-3} \text{ joule per second}}{3.14 \times 10^{-19} \text{ joule per photon}} = 3.2 \times 10^{15} \text{ photons per second}$$

This number is so large that the laser beam seems like a continuous flow of energy rather than a stream of little energy packets.



## 5-6 Each chemical element produces its own unique set of spectral lines

In 1814 the German master optician Joseph von Fraunhofer repeated the classic experiment of shining a beam of sunlight through a prism (see Figure 5-3). But this time Fraunhofer subjected the resulting rainbow-colored spectrum to intense magnification. To his surprise, he discovered that the solar spectrum contains hundreds of fine, dark lines, now called **spectral lines.** By contrast, if the light from a perfect blackbody were sent through a prism, it would produce a smooth, continuous spectrum with no dark lines. Fraunhofer counted more than 600 dark lines in the Sun's spectrum; today we know of more than 30,000. The photograph of the Sun's spectrum in Figure 5-13 shows hundreds of these spectral lines.

**Spectroscopy is the key to determining the chemical composition of planets and stars**

### Spectral Analysis

Half a century later, chemists discovered that they could produce spectral lines in the laboratory and use these spectral lines to analyze what kinds of atoms different substances are made of. Chemists had long known that many substances emit distinctive colors when sprinkled into a flame. To facilitate study of these colors, around 1857 the German chemist Robert Bunsen invented a gas burner (today called a Bunsen burner) that produces a clean flame with no color of its own. Bunsen's colleague, the Prussian-born physicist Gustav Kirchhoff, suggested that the colored light produced by substances in a flame might best be studied by passing the light through a prism (Figure 5-14). The two scientists promptly discovered that the spectrum from a flame consists of a pattern of thin, bright spectral lines against a dark background. The same kind of spectrum is produced by heated gases such as neon or argon.

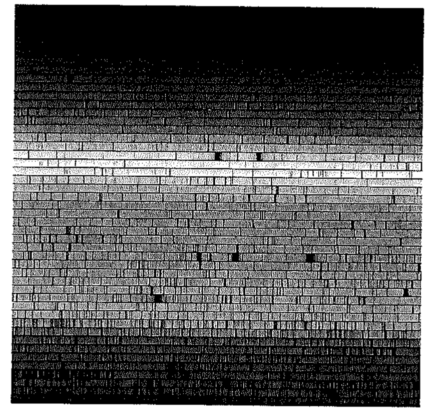


**Figure 5-13** R I V U X G

The Sun's Spectrum Numerous dark spectral lines are seen in this image of the Sun's spectrum. The spectrum is spread out so much that it had to be cut into segments to fit on this page. (N. A. Sharp, NOAO/NSO/Kitt Peak FTS/AURA/NSF)

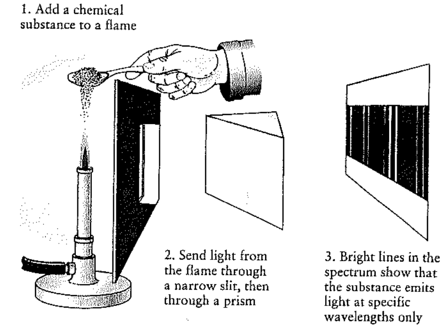



**Figure 5-14**

The Kirchhoff-Bunsen Experiment In the mid-1850s, Gustav Kirchhoff and Robert Bunsen discovered that when a chemical substance is heated and vaporized, the spectrum of the emitted light exhibits a series of bright spectral lines. They also found that each chemical element produces its own characteristic pattern of spectral lines. (In an actual laboratory experiment, lenses would be needed to focus the image of the slit onto the screen.)

Kirchhoff and Bunsen then found that each chemical element produces its own unique pattern of spectral lines. Thus was born in 1859 the technique of **spectral analysis,** the identification of chemical substances by their unique patterns of spectral lines.

A chemical **element** is a fundamental substance that cannot be broken down into more basic chemicals. Some examples are hydrogen, oxygen, carbon, iron, gold, and silver. After Kirchhoff and Bunsen had recorded the prominent spectral lines of all the then-known elements, they soon began to discover other spectral lines in the spectra of vaporized mineral samples. In this way they discovered elements whose presence had never before been suspected. In 1860, Kirchhoff and Bunsen found a new line in the blue portion of the spectrum of a sample of mineral water. After isolating the previously unknown element responsible for making the line, they named it cesium (from the Latin *caesium,* "gray-blue"). The next year, a new line in the red portion of the spectrum of a mineral sample led them to discover the element rubidium (Latin *rubidium,* "red").

Spectral analysis even allowed the discovery of new elements outside Earth. During the solar eclipse of 1868, astronomers found a new spectral line in light coming from the hot gases at the upper surface of the Sun while the main body of the Sun was hidden by the Moon. This line was attributed to a new element

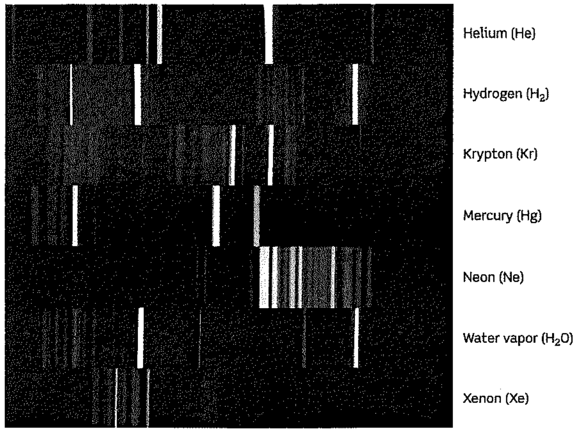



**Figure 5-15** R I V U X G

Various Spectra These photographs show the spectra of different types of gases as measured in a laboratory on Earth. Each type of gas has a unique spectrum that is the same wherever in the universe the gas is found. Water vapor ($H_2O$) is a compound whose molecules are made up of hydrogen and oxygen atoms; the hydrogen molecule ($H_2$) is made up of two hydrogen atoms. (Ted Kinsman/Science Photo Library)

that was named helium (from the Greek *helios,* "sun"). Helium was not discovered on Earth until 1895, when it was found in gases obtained from a uranium mineral.

A sample of an element contains only a single type of atom; carbon contains only carbon atoms, helium contains only helium atoms, and so on. Atoms of the same or different elements can combine to form **molecules**. For example, two hydrogen atoms (symbol H) can combine with an oxygen atom (symbol O) to form a water molecule (symbol $H_2O$). Substances like water whose molecules include atoms of different elements are called **compounds**. Just as each type of atom has its own unique spectrum, so does each type of molecule. Figure 5-15 shows the spectra of several types of atoms and molecules.

## Kirchhoff's Laws

The spectrum of the Sun, with its dark spectral lines superimposed on a bright background (see Figure 5-13), may seem to be unrelated to the spectra of bright lines against a dark background produced by substances in a flame (see Figure 5-14). But by the early 1860s, Kirchhoff's experiments had revealed a direct connection between these two types of spectra. His conclusions are summarized in three important statements about spectra that are today called **Kirchhoff's laws**. These laws, which are illustrated in Figure 5-16, are as follows:

Law 1 A hot opaque body, such as a perfect blackbody, or a hot, dense gas produces a **continuous spectrum**—a complete rainbow of colors without any spectral lines.

Law 2 A hot, transparent gas produces an **emission line spectrum**—a series of bright spectral lines against a dark background.

Law 3 A cool, transparent gas in front of a source of a continuous spectrum produces an **absorption line spectrum**—a series of dark spectral lines among the colors of the continuous spectrum. Furthermore, the dark lines in the absorption spectrum of a particular gas occur at exactly the *same* wavelengths as the bright lines in the emission spectrum of that same gas.

Kirchhoff's laws imply that if a beam of white light is passed through a gas, the atoms of the gas somehow extract light of very specific wavelengths from the white light. Hence, an observer who looks straight through the gas at the white-light source (the blackbody in Figure 5-16) will receive light whose spectrum has dark absorption lines superimposed on the continuous spectrum of the white light. The gas atoms then radiate light of precisely these same wavelengths in all directions. An observer at an oblique angle (that is, one who is not sighting directly through the cloud toward the blackbody) will receive only this light radiated by the gas cloud; the spectrum of this light is bright emission lines on a dark background.

**CAUTION!** Figure 5-16 shows that light can either pass through a cloud of gas or be absorbed by the gas. But there is also a third possibility: The light can simply bounce off the atoms or molecules that make up the gas, a phenomenon called **light scattering**. In other words, photons passing through a gas cloud can miss the gas atoms altogether, be swallowed whole by the atoms (absorption), or bounce off the atoms like billiard balls colliding (scattering). Box 5-4 describes how light scattering explains the blue color of the sky and the red color of sunsets.

Whether an emission line spectrum or an absorption line spectrum is observed from a gas cloud depends on the relative temperatures of the gas cloud and its background. Absorption lines are seen if the background is hotter than the gas, and emission lines are seen if the background is cooler.



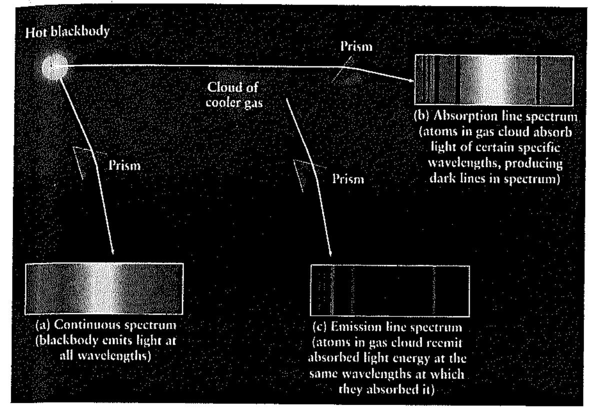



**Figure 5-16**

Continuous, Absorption Line, and Emission Line Spectra A hot, opaque body (like a blackbody) emits a continuous spectrum of light (spectrum *a*). If this light is passed through a cloud of a cooler gas, the cloud absorbs light of certain specific wavelengths, and the spectrum of light that passes directly through the cloud has dark absorption lines (spectrum *b*). The cloud does not retain all the light energy that it absorbs but radiates it outward in all directions. The spectrum of this reradiated light contains bright emission lines (spectrum *c*) with exactly the same wavelengths as the dark absorption lines in spectrum *b*. The specific wavelengths observed depend on the chemical composition of the cloud.

For example, if sodium is placed in the flame of a Bunsen burner in a darkened room, the flame will emit a characteristic orange-yellow glow. (This same glow is produced if we use ordinary table salt, which is a compound of sodium and chlorine.) If we pass the light from the flame through a prism, it displays an emission line spectrum with two closely spaced spectral lines at wavelengths of 588.99 and 589.59 nm, in the orange-yellow part of the spectrum. We now turn on a lightbulb whose filament is hotter than the flame and shine the bulb's white light through the flame. The spectrum of this light after it passes through the flame's sodium vapor is the continuous spectrum from the lightbulb, but with two closely spaced *dark* lines at 588.99 and 589.59 nm. Thus, the chemical composition of the gas is revealed by either bright emission lines or dark absorption lines.

## Spectroscopy

**Spectroscopy** is the systematic study of spectra and spectral lines. Spectral lines are tremendously important in astronomy, because they provide reliable evidence about the chemical composition of distant objects. As an example, the spectrum of the Sun shown in Figure 5-13 is an absorption line spectrum. The continuous spectrum comes from the hot surface of the Sun, which acts like a blackbody. The dark absorption lines are caused by this light passing through a cooler gas; this gas is the atmosphere that surrounds the Sun. Therefore, by identifying the spectral lines present in the solar spectrum, we can determine the chemical composition of the Sun's atmosphere.

Figure 5-17 shows both a portion of the Sun's absorption line spectrum and the emission line spectrum of iron vapor over the same wavelength range. This pattern of bright spectral lines in the lower spectrum is iron's own distinctive "fingerprint," which no other substance can imitate. Because some absorption lines in the Sun's spectrum coincide with the iron lines, some vaporized iron must exist in the Sun's atmosphere.

Spectroscopy can also help us analyze gas clouds in space, such as the nebula surrounding the star cluster NGC 346 shown

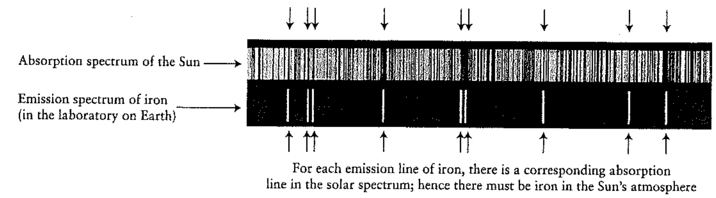



**Figure 5-17** R I V U X G

Iron in the Sun The upper part of this figure is a portion of the Sun's spectrum at violet wavelengths, showing numerous dark absorption lines. The lower part of the figure is a corresponding portion of the emission line spectrum of vaporized iron. The iron lines coincide with some of the solar lines, which proves that there is some iron (albeit a relatively small amount) in the Sun's atmosphere. (Carnegie Observatories)

BOX 5-4 Astronomy Down to Earth

## Light Scattering

Light scattering is the process whereby photons bounce off particles in their path. These particles can be atoms, molecules, or clumps of molecules. You are reading these words using photons from the Sun or a lamp that bounced off the page—that is, were scattered by the particles that make up the page.

An important fact about light scattering is that very small particles—ones that are smaller than a wavelength of visible light—are quite effective at scattering short-wavelength photons of blue light, but less effective at scattering long-wavelength photons of red light. This fact explains a number of phenomena that you can see here on Earth.

The light that comes from the daytime sky is sunlight that has been scattered by the molecules that make up our atmosphere (see part *a* of the accompanying figure). Air molecules are less than 1 nm across, far smaller than the wavelength of visible light, so they scatter blue light more than red light—which is why the sky looks blue. Smoke particles are also quite small, which explains why the smoke from a cigarette or a fire has a bluish color.

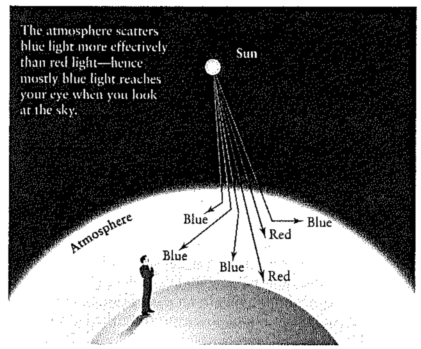
Distant mountains often appear blue thanks to sunlight being scattered from the atmosphere between the mountains and your eyes. (The Blue Ridge Mountains, which extend from Pennsylvania to Georgia, and Australia's Blue Mountains derive their names from this effect.) Sunglasses often have a red or orange tint, which blocks out blue light. This cuts down on the amount of scattered light from the sky reaching your eyes and allows you to see distant objects more clearly.

Light scattering also explains why sunsets are red. The light from the Sun contains photons of all visible wavelengths, but as this light passes through our atmosphere the blue photons are scattered away from the straight-line path from the Sun to your eye. Red photons undergo relatively little scattering, so the Sun always looks a bit redder than it really is. When you look toward the setting sun, the sunlight that reaches your eye has had to pass through a relatively thick layer of atmosphere (part *b* of the accompanying figure). Hence, a large fraction of the blue light from the Sun has been scattered, and the Sun appears quite red.

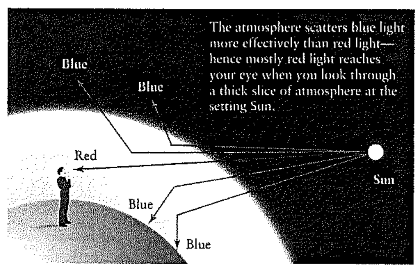
The same effect also applies to sunrises, but sunrises seldom look as red as sunsets do. The reason is that dust is lifted into the atmosphere during the day by the wind (which is typically stronger in the daytime than at night), and dust particles in the atmosphere help to scatter even more blue light.

If the small particles that scatter light are sufficiently concentrated, there will be almost as much scattering of red light as of blue light, and the scattered light will appear white. This explains the white color of clouds, fog, and haze, in which the scattering particles are ice crystals or water droplets. Whole milk looks white because of light scattering from tiny fat globules; nonfat milk has only a very few of these globules and so has a slight bluish cast.

Light scattering has many applications to astronomy. For example, it explains why very distant stars in our Galaxy appear surprisingly red. The reason is that there are tiny dust particles in the space between the stars, and this dust scatters blue photons. By studying how much scattering takes place, astronomers have learned about the tenuous material that fills interstellar space.




(a) Why the sky looks blue




(b) Why the setting Sun looks red

**Figure 5-18** RI**V**UXG

Analyzing the Composition of a Distant Nebula The glowing gas cloud in this Hubble Space Telescope image lies 210,000 light-years away in the constellation Tucana (the Toucan). Hot stars within the nebula emit high-energy, ultraviolet photons, which are absorbed by the surrounding gas and heat the gas to high temperature. This heated gas produces light with an emission line spectrum. The particular wavelength of red light emitted by the nebula is 656 nm, characteristic of hydrogen gas. (NASA, ESA, and A. Nota, STScI/ESA)

in Figure 5-18. Such glowing clouds have emission line spectra, because we see them against the black background of space. The particular shade of red that dominates the color of this nebula is due to an emission line at a wavelength near 656 nm. This is one of the characteristic wavelengths emitted by hydrogen gas, so we can conclude that this nebula contains hydrogen. More detailed analyses of this kind show that hydrogen is the most common element in gaseous nebulae, and indeed in the universe as a whole. The spectra of other nebulae, such as the Ring Nebula shown in the image that opens this chapter, also reveal the presence of nitrogen, oxygen, helium, and other gases.

What is truly remarkable about spectroscopy is that it can determine chemical composition at any distance. The 656-nm red light produced by a sample of heated hydrogen gas on Earth (the bright red line in the hydrogen spectrum in Figure 5-15) is the same as that observed coming from the nebula shown in Figure 5-18, located about 210,000 light-years away. By using the basic principles outlined by Kirchhoff, astronomers have the tools to make chemical assays of objects that are almost inconceivably distant. Throughout this book we will see many examples of how astronomers use Kirchhoff's laws to determine the nature of celestial objects.

To make full use of Kirchhoff's laws, it is helpful to understand why they work. Why does an atom absorb light of only particular wavelengths? And why does it then emit light of only these same wavelengths? Maxwell's theory of electromagnetism (see Section 5-2) could not answer these questions. The answers did not come until early in the twentieth century, when scientists began to discover the structure and properties of atoms.

## 5-7 An atom consists of a small, dense nucleus surrounded by electrons

**To decode the information in the light from immense objects like stars and galaxies, we must understand the structure of atoms**

The first important clue about the internal structure of atoms came from an experiment conducted in 1910 by Ernest Rutherford, a gifted physicist from New Zealand. Rutherford and his colleagues at the University of Manchester in England had been investigating the recently discovered phenomenon of radioactivity. Certain radioactive elements, such as uranium and radium, were known to emit particles of various types. One type, the alpha particle, has about the same mass as a helium atom and is emitted from some radioactive substances with considerable speed.

In one series of experiments, Rutherford and his colleagues were using alpha particles as projectiles to probe the structure of solid matter. They directed a beam of these particles at a thin sheet of metal (Figure 5-19). Almost all the alpha particles passed through the metal sheet with little or no deflection from their straight-line paths. To the surprise of the experimenters, however, an occasional alpha particle bounced back from the metal sheet as though it had struck something quite dense. Rutherford later remarked, "It was almost as incredible as

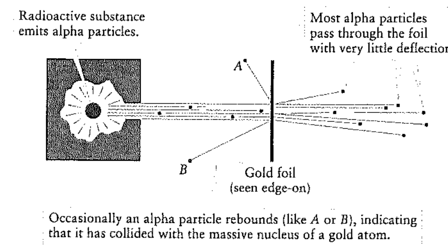



**Figure 5-19**

Rutherford's Experiment Alpha particles from a radioactive source are directed at a thin metal foil. This experiment provided the first evidence that the nuclei of atoms are relatively massive and compact.

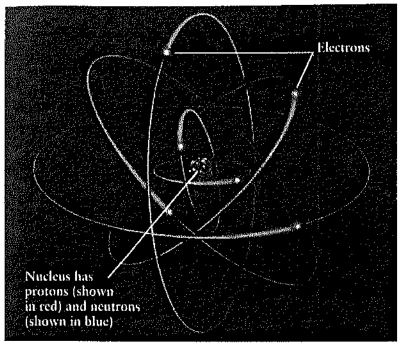



**Figure 5-20**

Rutherford's Model of the Atom In this model, electrons orbit the atom's nucleus, which contains most of the atom's mass. The nucleus contains two types of particles, protons and neutrons.

if you fired a fifteen-inch shell at a piece of tissue paper and it came back and hit you."

Rutherford concluded from this experiment that most of the mass of an atom is concentrated in a compact, massive lump of matter that occupies only a small part of the atom's volume. Most of the alpha particles pass freely through the nearly empty space that makes up most of the atom, but a few particles happen to strike the dense mass at the center of the atom and bounce back.

### The Nucleus of an Atom

Rutherford proposed a new model for the structure of an atom, shown in Figure 5-20. According to this model, a massive, positively charged **nucleus** at the center of the atom is orbited by tiny, negatively charged electrons. Rutherford concluded that at least 99.98% of the mass of an atom must be concentrated in its nucleus, whose diameter is only about $10^{-14}$ m. (The diameter of a typical atom is far larger, about $10^{-10}$ m.)

ANALOGY To appreciate just how tiny the nucleus is, imagine expanding an atom by a factor of $10^{12}$ to a diameter of 100 meters, about the length of a football field. To this scale, the nucleus would be just a centimeter across—no larger than your thumbnail.

We know today that the nucleus of an atom contains two types of particles, **protons** and **neutrons**. A proton has a positive electric charge, equal and opposite to that of an electron. As its name suggests, a neutron has no electric charge—it is electrically neutral. As an example, an alpha particle (such as those Rutherford's team used) is actually a nucleus of the helium atom, with two protons and two neutrons. Protons and neutrons are held together in a nucleus by the so-called strong nuclear force, whose great strength overcomes the electric repulsion between the positively charged protons. A proton and a neutron have almost the same mass, $1.7 \times 10^{-27}$ kg, and each has about 2000 times as much mass as an electron ($9.1 \times 10^{-31}$ kg). In an ordinary atom there are as many positive protons as there are negative electrons, so the atom has no net electric charge. Because the mass of the electron is so small, the mass of an atom is not much greater than the mass of its nucleus. That is why an alpha particle has nearly the same mass as an atom of helium.

While the solar system is held together by gravitational forces, atoms are held together by electrical forces. Opposites attract: The negative charges on the orbiting electrons are attracted to the positive charges on the protons in the nucleus. Box 5-5 on page 118 describes more about the connection between the structure of atoms and the chemical and physical properties of substances made of those atoms.

Rutherford's experiments clarified the structure of the atom, but they did not explain how these tiny particles within the atom give rise to spectral lines. The task of reconciling Rutherford's atomic model with Kirchhoff's laws of spectral analysis was undertaken by the young Danish physicist Niels Bohr, who joined Rutherford's group at Manchester in 1912.

## 5-8 Spectral lines are produced when an electron jumps from one energy level to another within an atom

WEB LINK 5.8

Niels Bohr began his study of the connection between atomic spectra and atomic structure by trying to understand the structure of hydrogen, the simplest and lightest of the elements. (As we discussed in Section 5-6, hydrogen is also the most common element in the universe.) When Bohr was done, he had not only found a way to explain this atom's spectrum but had also found a justification for Kirchhoff's laws in terms of atomic physics.

**Niels Bohr explained spectral lines with a radical new model of the atom**

### Hydrogen and the Balmer Series

The most common type of hydrogen atom consists of a single electron and a single proton. Hydrogen atoms have a simple visible-light spectrum consisting of a pattern of lines that begins at a wavelength of 656.3 nm and ends at 364.6 nm. The first spectral line is called $H_\alpha$ (H-alpha), the second spectral line is called $H_\beta$ (H-beta), the third is $H_\gamma$ (H-gamma), and so forth. (These are the bright lines in the spectrum of hydrogen shown in Figure 5-15. The fainter lines between these appear when hydrogen atoms form into molecules.) The closer you get to the short-wavelength end of the spectrum at 364.6 nm, the more spectral lines you see.

The regularity in this spectral pattern was described mathematically in 1885 by Johann Jakob Balmer, a Swiss schoolteacher. The spectral lines of hydrogen at visible wavelengths are today called **Balmer lines**, and the entire pattern from $H_\alpha$ onward is called the **Balmer series**. Eight Balmer lines are seen in the spectrum of the star shown in Figure 5-21. Stars in general, including the Sun, have Balmer absorption lines in their spectra, which shows they have atmospheres that contain hydrogen.

Using trial and error, Balmer discovered a formula from which the wavelengths (λ) of hydrogen's spectral lines can be calculated. Balmer's formula is usually written

$$\frac{1}{\lambda} = R\left(\frac{1}{4} - \frac{1}{n^2}\right)$$

In this formula $R$ is the *Rydberg constant* ($R = 1.097 \times 10^7$ m$^{-1}$), named in honor of the Swedish spectroscopist Johannes Rydberg, and $n$ can be any integer (whole number) greater than 2. To get the wavelength $\lambda_\alpha$ of the spectral line $H_\alpha$, you first put $n = 3$ into Balmer's formula:

$$\frac{1}{\lambda_\alpha} = (1.097 \times 10^7 \text{ m}^{-1})\left(\frac{1}{4} - \frac{1}{3^2}\right) = 1.524 \times 10^6 \text{ m}^{-1}$$

Then take the reciprocal:

$$\lambda_\alpha = \frac{1}{1.524 \times 10^6 \text{ m}^{-1}} = 6.563 \times 10^{-7} \text{ m} = 656.3 \text{ nm}$$

To get the wavelength of $H_\beta$, use $n = 4$, and to get the wavelength of $H_\gamma$, use $n = 5$. If you use $n = \infty$ (the symbol ∞ stands for infinity), you get the short-wavelength end of the hydrogen spectrum at 364.6 nm. (Note that 1 divided by infinity equals zero.)

### Bohr's Model of Hydrogen

Bohr realized that to fully understand the structure of the hydrogen atom, he had to be able to derive Balmer's formula using the laws of physics. He first made the rather wild assumption that the electron in a hydrogen atom can orbit the nucleus only in certain specific orbits. (This idea was a significant break with the ideas of Newton, in whose mechanics any orbit should be possible.) Figure 5-22 shows the four smallest of these **Bohr orbits**, labeled by the numbers $n = 1$, $n = 2$, $n = 3$, and so on.

Although confined to one of these allowed orbits while circling the nucleus, an electron can jump from one Bohr orbit to another. For an electron to jump, the hydrogen atom must gain or lose a specific amount of energy. The atom must absorb energy for the electron to go from an inner to an outer orbit; the atom must release energy for the electron to go from an outer to an inner orbit. As an example, Figure 5-23 shows an electron jumping between the $n = 2$ and $n = 3$ orbits of a hydrogen atom as the atom absorbs or emits an $H_\alpha$ photon.

When the electron jumps from one orbit to another, the energy of the photon that is emitted or absorbed equals the difference in energy between these two orbits. This energy difference, and hence the photon energy, is the same whether the jump is from a low orbit to a high orbit (Figure 5-23*a*) or from the high orbit back to the low one (Figure 5-23*b*). According to Planck and Einstein, if two photons have the same energy $E$, the relationship $E = hc/\lambda$ tells us that they must also have the same wavelength λ. It follows that if an atom can emit photons of a given energy and wavelength, it can also absorb photons of precisely the same energy and wavelength. Thus, Bohr's picture explains Kirchhoff's observation that atoms emit and absorb the same wavelengths of light.

The Bohr picture also helps us visualize what happens to produce an emission line spectrum. When a gas is heated, its atoms move around rapidly and can collide forcefully with each other. These energetic collisions excite the atoms' electrons into high orbits. The electrons then cascade back down to the innermost

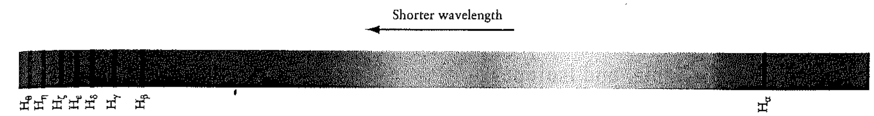



**Figure 5-21** R I V U X G

Balmer Lines in the Spectrum of a Star This portion of the spectrum of the star Vega in the constellation Lyra (the Harp) shows eight Balmer lines, from $H_\alpha$ at 656.3 nm through $H_\theta$ (H-theta) at 388.9 nm. The series converges at 364.6 nm, slightly to the left of $H_\theta$. Parts of this image were made using ultraviolet radiation, which is indicated by the highlighted U in the wavelength tab. (NOAO)

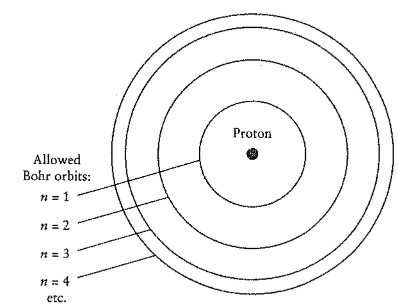



**Figure 5-22**

The Bohr Model of the Hydrogen Atom In this model, an electron circles the hydrogen nucleus (a proton) only in allowed orbits $n = 1, 2, 3$, and so forth. The first four Bohr orbits are shown here. This figure is *not* drawn to scale; in the Bohr model, the $n = 2, 3$, and 4 orbits are respectively 4, 9, and 16 times larger than the $n = 1$ orbit.

BOX 5-5 Tools of the Astronomer's Trade

## Atoms, the Periodic Table, and Isotopes

WEB LINK 5.6

Each different chemical element is made of a specific type of atom. Each specific atom has a characteristic number of protons in its nucleus. For example, a hydrogen atom has 1 proton in its nucleus, an oxygen atom has 8 protons in its nucleus, and so on.

The number of protons in an atom's nucleus is the **atomic number** for that particular element. The chemical elements are most conveniently listed in the form of a **periodic table** (shown in the figure). Elements are arranged in the periodic table in order of increasing atomic number. With only a few exceptions, this sequence also corresponds to increasing average mass of the atoms of the elements. Thus, hydrogen (symbol H), with atomic number 1, is the lightest element. Iron (symbol Fe) has atomic number 26 and is a relatively heavy element.

All the elements listed in a single vertical column of the periodic table have similar chemical properties. For example, the elements in the far right column are all gases under the conditions of temperature and pressure found at Earth's surface, and they are all very reluctant to react chemically with other elements.

In addition to nearly 100 naturally occurring elements, the periodic table includes a number of artificially produced elements. Most of these elements are heavier than uranium (symbol U) and are highly radioactive, which means that they decay into lighter elements within a short time of being created in laboratory experiments. Scientists have succeeded in creating only a few atoms of elements 104 and above.

The number of protons in the nucleus of an atom determines which element that atom is. Nevertheless, the same element may have different numbers of neutrons in its nucleus. For example, oxygen (O) has atomic number 8, so every oxygen nucleus has exactly 8 protons. But oxygen nuclei can have 8, 9, or 10 neutrons. These three slightly different kinds of oxygen are called **isotopes**. The isotope with 8 neutrons is by far the most abundant variety. It is written as $^{16}O$, or oxygen-16. The rarer isotopes with 9 and 10 neutrons are designated as $^{17}O$ and $^{18}O$, respectively.

The superscript that precedes the chemical symbol for an element equals the total number of protons and neutrons in a nucleus of that particular isotope. For example, a nucleus of the most common isotope of iron, $^{56}Fe$ or iron-56, contains a total of 56 protons and neutrons. From the periodic table, the atomic number of iron is 26, so every iron atom has 26 protons in its nucleus. Therefore, the number of neutrons in an iron-56 nucleus is $56 - 26 = 30$. (Most nuclei have more neutrons than protons, especially in the case of the heaviest elements.)

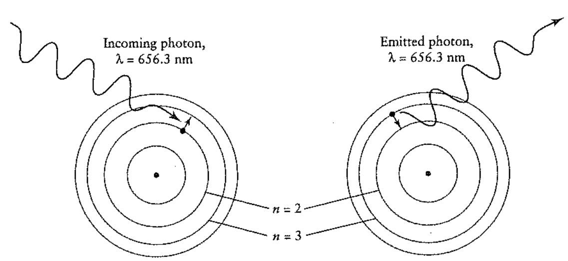
It is extremely difficult to distinguish chemically between the various isotopes of a particular element. Ordinary chemical reactions involve only the electrons that orbit the atom, never the neutrons buried in its nucleus. But there are small differences in the wavelengths of the spectral lines for different isotopes of the same element. For example, the spectral line wavelengths of the hydrogen isotope $^{2}H$ are about 0.03% greater than the wavelengths for the most common hydrogen isotope, $^{1}H$. Thus, different isotopes can be distinguished by careful spectroscopic analysis.

Isotopes are important in astronomy for a variety of reasons. By measuring the relative amounts of different isotopes of a given element in a Moon rock or meteorite, the age of that sample can be determined. The mixture of isotopes left behind when a star explodes into a supernova (see Section 1-3) tells astronomers about the processes that led to the explosion. And knowing the properties of different isotopes of hydrogen and helium is crucial to understanding the nuclear reactions that make the Sun shine. Look for these and other applications of the idea of isotopes in later chapters.

Periodic Table of the Elements

| | | | | | | | | | | | | | | | | | |
|---|---|---|---|---|---|---|---|---|---|---|---|---|---|---|---|---|---|
| 1 H Hydrogen | | | | | | | | | | | | | | | | | 2 He Helium |
| 3 Li Lithium | 4 Be Beryllium | | | | | | | | | | | 5 B Boron | 6 C Carbon | 7 N Nitrogen | 8 O Oxygen | 9 F Fluorine | 10 Ne Neon |
| 11 Na Sodium | 12 Mg Magnesium | | | | | | | | | | | 13 Al Aluminum | 14 Si Silicon | 15 P Phosphorus | 16 S Sulfur | 17 Cl Chlorine | 18 Ar Argon |
| 19 K Potassium | 20 Ca Calcium | 21 Sc Scandium | 22 Ti Titanium | 23 V Vanadium | 24 Cr Chromium | 25 Mn Manganese | 26 Fe Iron | 27 Co Cobalt | 28 Ni Nickel | 29 Cu Copper | 30 Zn Zinc | 31 Ga Gallium | 32 Ge Germanium | 33 As Arsenic | 34 Se Selenium | 35 Br Bromine | 36 Kr Krypton |
| 37 Rb Rubidium | 38 Sr Strontium | 39 Y Yttrium | 40 Zr Zirconium | 41 Nb Niobium | 42 Mo Molybdenum | 43 Tc Technetium | 44 Ru Ruthenium | 45 Rh Rhodium | 46 Pd Palladium | 47 Ag Silver | 48 Cd Cadmium | 49 In Indium | 50 Sn Tin | 51 Sb Antimony | 52 Te Tellurium | 53 I Iodine | 54 Xe Xenon |
| 55 Cs Cesium | 56 Ba Barium | 57 La Lanthanum | 72 Hf Hafnium | 73 Ta Tantalum | 74 W Tungsten | 75 Re Rhenium | 76 Os Osmium | 77 Ir Iridium | 78 Pt Platinum | 79 Au Gold | 80 Hg Mercury | 81 Tl Thallium | 82 Pb Lead | 83 Bi Bismuth | 84 Po Polonium | 85 At Astatine | 86 Rn Radon |
| 87 Fr Francium | 88 Ra Radium | 89 Ac Actinium | 104 Rf Rutherfordium | 105 Db Dubnium | 106 Sg Seaborgium | 107 Bh Bohrium | 108 Hs Hassium | 109 Mt Meitnerium | 110 Ds Darmstadtium | 111 Rg Roentgenium | | | | | | | |

| | | | | | | | | | | | | | |
|---|---|---|---|---|---|---|---|---|---|---|---|---|---|
| 58 Ce Cerium | 59 Pr Praseodymium | 60 Nd Neodymium | 61 Pm Promethium | 62 Sm Samarium | 63 Eu Europium | 64 Gd Gadolinium | 65 Tb Terbium | 66 Dy Dysprosium | 67 Ho Holmium | 68 Er Erbium | 69 Tm Thulium | 70 Yb Ytterbium | 71 Lu Lutetium |
| 90 Th Thorium | 91 Pa Protactinium | 92 U Uranium | 93 Np Neptunium | 94 Pu Plutonium | 95 Am Americium | 96 Cm Curium | 97 Bk Berkelium | 98 Cf Californium | 99 Es Einsteinium | 100 Fm Fermium | 101 Md Mendelevium | 102 No Nobelium | 103 Lr Lawrencium |




(a) Atom absorbs a 656.3-nm photon; absorbed energy causes electron to jump from the $n = 2$ orbit up the $n = 3$ orbit

(b) Electron falls from the $n = 3$ orbit to the $n = 2$ orbit; energy lost by atom goes into emitting a 656.3-nm photon

ANIMATION 5.2

**Figure 5-23**

The Absorption and Emission of an $H_\alpha$ Photon This schematic diagram, drawn according to the Bohr model, shows what happens when a hydrogen atom (a) absorbs or (b) emits a photon whose wavelength is 656.3 nm.

possible orbit, emitting photons whose energies are equal to the energy differences between different Bohr orbits. In this fashion, a hot gas produces an emission line spectrum with a variety of different wavelengths.

To produce an absorption line spectrum, begin with a relatively cool gas, so that the electrons in most of the atoms are in inner, low-energy orbits. If a beam of light with a continuous spectrum is shone through the gas, most wavelengths will pass through undisturbed. Only those photons will be absorbed whose energies are just right to excite an electron to an allowed outer orbit. Hence, only certain wavelengths will be absorbed, and dark lines will appear in the spectrum at those wavelengths.

Using his picture of allowed orbits and the formula $E = hc/\lambda$, Bohr was able to prove mathematically that the wavelength λ of the photon emitted or absorbed as an electron jumps between an inner orbit $N$ and an outer orbit $n$ is

**Bohr formula for hydrogen wavelengths**

$$\frac{1}{\lambda} = R\left(\frac{1}{N^2} - \frac{1}{n^2}\right)$$

$N$ = number of inner orbit

$n$ = number of outer orbit

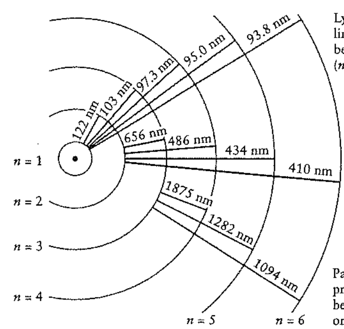
$R$ = Rydberg constant = $1.097 \times 10^7\ m^{-1}$

λ = wavelength (in meters) of emitted or absorbed photon

If Bohr let $N = 2$ in this formula, he got back the formula that Balmer discovered by trial and error. Hence, Bohr deduced the meaning of the Balmer series: All the Balmer lines are produced by electrons jumping between the second Bohr orbit

## Figure 5-24

Electron Transitions in the Hydrogen Atom This diagram shows the photon wavelengths associated with different electron transitions in hydrogen. In each case, the same wavelength occurs whether a photon is emitted (when the electron drops from a high orbit to a low one) or absorbed (when the electron jumps from a low orbit to a high one). The orbits are not shown to scale.




($N = 2$) and higher orbits ($n = 3, 4, 5$, and so on). Remarkably, as part of his calculation Bohr was able to *derive* the value of the Rydberg constant in terms of Planck's constant, the speed of light, and the mass and electric charge of the electron. This gave particular credence to his radical model.

Bohr's formula also correctly predicts the wavelengths of other series of spectral lines that occur at nonvisible wavelengths (Figure 5-24). Using $N = 1$ gives the **Lyman series,** which is entirely in the ultraviolet. All the spectral lines in this series involve electron transitions between the lowest Bohr orbit and all higher orbits ($n = 2, 3, 4$, and so on). This pattern of spectral lines begins with $L_\alpha$ (Lyman alpha) at 122 nm and converges on $L_\infty$ at 91.2 nm. Using $N = 3$ gives a series of infrared wavelengths called the **Paschen series.** This series, which involves transitions between the third Bohr orbit and all higher orbits, begins with $P_\alpha$ (Paschen alpha) at 1875 nm and converges on $P_\infty$ at 822 nm. Additional series exist at still longer wavelengths.

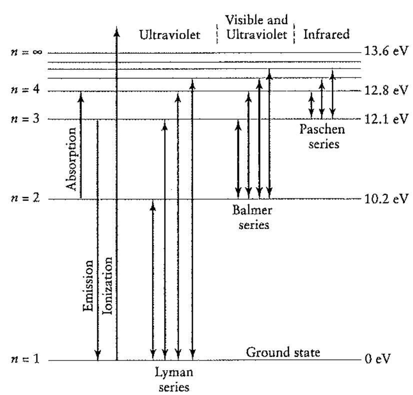
### Atomic Energy Levels

Today's view of the atom owes much to the Bohr model, but is different in certain ways. The modern picture is based on **quantum mechanics,** a branch of physics developed during the 1920s that deals with photons and subatomic particles. As a result of this work, physicists no longer picture electrons as moving in specific orbits about the nucleus. Instead, electrons are now known to have both particle and wave properties and are said to occupy only certain **energy levels** in the atom.

An extremely useful way of displaying the structure of an atom is with an **energy-level diagram.** Figure 5-25 shows such a diagram for hydrogen. The lowest energy level, called the **ground state,** corresponds to the $n = 1$ Bohr orbit. Higher energy levels, called **excited states,** correspond to successively larger Bohr orbits.

An electron can jump from the ground state up to the $n = 2$ level if the atom absorbs a Lyman-alpha photon with a wavelength of 122 nm. Such a photon has energy $E = hc/\lambda = 10.2$ eV (electron volts; see Section 5-5). That's why the energy level of $n = 2$ is shown in Figure 5-25 as having an energy 10.2 eV above that of the ground state (which is usually assigned a value of 0 eV). Similarly, the $n = 3$ level is 12.1 eV above the ground state, and so forth. Electrons can make transitions to higher energy levels by absorbing a photon or in a collision between atoms; they can make transitions to lower energy levels by emitting a photon.

On the energy-level diagram for hydrogen, the $n = \infty$ level has an energy of 13.6 eV. (This corresponds to an infinitely large orbit in the Bohr model.) If the electron is initially in the ground state and the atom absorbs a photon of any energy greater than 13.6 eV, the electron will be removed completely from the atom. This process is called **ionization.** A 13.6-eV photon has a wavelength of 91.2 nm, equal to the shortest wavelength in the ultraviolet Lyman series ($L_\infty$). So any photon with a wavelength of 91.2 nm or less can ionize hydrogen. (The Planck formula $E = hc/\lambda$ tells us that the higher the photon energy, the shorter the wavelength.)




## Figure 5-25

Energy-Level Diagram of Hydrogen A convenient way to display the structure of the hydrogen atom is in a diagram like this, which shows the allowed energy levels. The diagram shows a number of possible electron jumps, or transitions, between energy levels. An upward transition occurs when the atom absorbs a photon; a downward transition occurs when the atom emits a photon. (Compare with Figure 5-24.)



As an example, the gaseous nebula shown in Figure 5-18 surrounds a cluster of hot stars which produce copious amounts of ultraviolet photons with wavelengths less than 91.2 nm. Hydrogen atoms in the nebula that absorb these photons become ionized and lose their electrons. When the electrons recombine with the nuclei, they cascade down the energy levels to the ground state and emit visible light in the process. This process is what makes the nebula glow.

### The Spectra of Other Elements

The same basic principles that explain the hydrogen spectrum also apply to the atoms of other elements. Electrons in each kind of atom can be only in certain energy levels, so only photons of certain wavelengths can be emitted or absorbed. Because each kind of atom has its own unique arrangement of electron levels, the pattern of spectral lines is likewise unique to that particular type of atom (see the photograph that opens this chapter). These patterns are in general much more complicated than for the hydrogen atom. Hence, there is no simple relationship analogous to the Bohr formula that applies to the spectra of all atoms.

The idea of energy levels explains the emission line spectra and absorption line spectra of gases. But what about the continuous spectra produced by dense objects like the filament of a lightbulb or the coils of a toaster? These objects are made of atoms, so why don't they emit light with an emission line spectrum characteristic of the particular atoms of which they are made?

The reason is directly related to the difference between a gas on the one hand and a liquid or solid on the other. In a gas, atoms are widely separated and can emit photons without interference from other atoms. But in a liquid or a solid, atoms are so close that they almost touch, and thus these atoms interact strongly with each other. These interactions interfere with the process of emitting photons. As a result, the pattern of distinctive bright spectral lines that the atoms would emit in isolation becomes "smeared out" into a continuous spectrum.

ANALOGY Think of atoms as being like tuning forks. If you strike a single tuning fork, it produces a sound wave with a single clear frequency and wavelength, just as an isolated atom emits light of definite wavelengths. But if you shake a box packed full of tuning forks, you will hear a clanging noise that is a mixture of sounds of all different frequencies and wavelengths. This is directly analogous to the continuous spectrum of light emitted by a dense object with closely packed atoms.

With the work of such people as Planck, Einstein, Rutherford, and Bohr, the interchange between astronomy and physics came full circle. Modern physics was born when Newton set out to understand the motions of the planets. Two and a half centuries later, physicists in their laboratories probed the properties of light and the structure of atoms. Their labors had immediate applications in astronomy. Armed with this new understanding of light and matter, astronomers were able to probe in detail the chemical and physical properties of planets, stars, and galaxies.

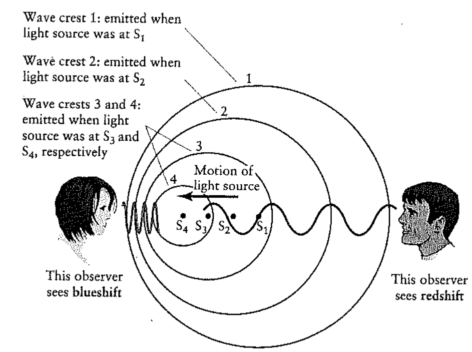
## 5-9 The wavelength of a spectral line is affected by the relative motion between the source and the observer

In addition to telling us about temperature and chemical composition, the spectrum of a planet, star, or galaxy can also reveal something about that object's motion through space. This idea dates from 1842, when Christian Doppler, a professor of mathematics in Prague, pointed out that the observed wavelength of light must be affected by motion.

**The Doppler effect makes it possible to tell whether astronomical objects are moving toward us or away from us**

### The Doppler Effect

In Figure 5-26 a light source is moving from right to left; the circles represent the crests of waves emitted from the moving source at various positions. Each successive wave crest is emitted from a position slightly closer to the observer on the left, so she sees a shorter wavelength—the distance from one crest to the next—than she would if the source were stationary. All the lines in the spectrum




## Figure 5-26

The Doppler Effect The wavelength of light is affected by motion between the light source and an observer. The light source shown here is moving, so wave crests 1, 2, etc., emitted when the source was at points $S_1$, $S_2$, etc., are crowded together in front of the source but are spread out behind it. Consequently, wavelengths are shortened (blueshifted) if the source is moving toward the observer and lengthened (redshifted) if the source is moving away from the observer. Motion perpendicular to an observer's line of sight does not affect wavelength.

**BOX 5-6** Tools of the Astronomer's Trade

## Applications of the Doppler Effect

Doppler's formula relates the radial velocity of an astronomical object to the wavelength shift of its spectral lines. Here are two examples that show how to use this remarkably powerful formula.

EXAMPLE: As measured in the laboratory, the prominent $H_\alpha$ spectral line of hydrogen has a wavelength $\lambda_0 = 656.285$ nm. But in the spectrum of the star Vega (Figure 5-21), this line has a wavelength $\lambda = 656.255$ nm. What can we conclude about the motion of Vega?

Situation: Our goal is to use the ideas of the Doppler effect to find the velocity of Vega toward or away from Earth.

Tools: We use the Doppler shift formula, $\lambda/\lambda_0 = v/c$, to determine Vega's velocity $v$.

Answer: The wavelength shift is

$$\Delta\lambda = \lambda - \lambda_0 = 656.255 \text{ nm} - 656.285 \text{ nm} = -0.030 \text{ nm}$$

The negative value means that we see the light from Vega shifted to shorter wavelengths—that is, there is a blueshift. (Note that the shift is very tiny and can be measured only using specialized equipment.) From the Doppler shift formula, the star's radial velocity is

$$v = c\frac{\Delta\lambda}{\lambda_0} = (3.00 \times 10^5 \text{ km/s})\left(\frac{-0.030 \text{ nm}}{656.285 \text{ nm}}\right) = -14 \text{ km/s}$$

Review: The minus sign indicates that Vega is coming toward us at 14 km/s. The star may also be moving perpendicular to the line from Earth to Vega, but such motion produces no Doppler shift.

By plotting the motions of stars such as Vega toward and away from us, astronomers have been able to learn how the Milky Way Galaxy (of which our Sun is a part) is rotating. From this knowledge, and aided by Newton's universal law of gravitation (see Section 4-7), they have made the surprising discovery that the Milky Way contains roughly 10 times more matter than had once been thought! The nature of this unseen *dark matter* is still a subject of debate.

EXAMPLE: In the radio region of the electromagnetic spectrum, hydrogen atoms emit and absorb photons with a wavelength of 21.12 cm, giving rise to a spectral feature commonly called the *21-centimeter line*. The galaxy NGC 3840 in the constellation Leo (the Lion) is receding from us at a speed of 7370 km/s, or about 2.5% of the speed of light. At what wavelength do we expect to detect the 21-cm line from this galaxy?

Situation: Given the velocity of NGC 3840 away from us, our goal is to find the wavelength as measured on Earth of the 21-centimeter line from this galaxy.

Tools: We use the Doppler shift formula to calculate the wavelength shift $\Delta\lambda$, then use this to find the wavelength $\lambda$ measured on Earth.

Answer: The wavelength shift is

$$\Delta\lambda = \lambda_0\left(\frac{v}{c}\right) = (21.12 \text{ cm})\left(\frac{7370 \text{ km/s}}{3.00 \times 10^5 \text{ km/s}}\right) = 0.52 \text{ cm}$$

Therefore, we will detect the 21-cm line of hydrogen from this galaxy at a wavelength of

$$\lambda = \lambda_0 + \Delta\lambda = 21.12 \text{ cm} + 0.52 \text{ cm} = 21.64 \text{ cm}$$

Review: The 21-cm line has been redshifted to a longer wavelength because the galaxy is receding from us. In fact, most galaxies are receding from us. This observation is one of the key pieces of evidence that the universe is expanding, and has been doing so since the Big Bang that took place almost 14 billion years ago.

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of an approaching source are shifted toward the short-wavelength (blue) end of the spectrum. This phenomenon is called a **blueshift.**

The source is receding from the observer on the right in Figure 5-26. The wave crests that reach him are stretched apart, so that he sees a longer wavelength than he would if the source were stationary. All the lines in the spectrum of a receding source are shifted toward the longer-wavelength (red) end of the spectrum, producing a **redshift.** In general, the effect of relative motion on wavelength is called the **Doppler effect.** Police radar guns use the Doppler effect to check for cars exceeding the speed limit: The radar gun sends a radio wave toward the car, and measures the wavelength shift of the reflected wave (and thus the speed of the car).

ANALOGY You have probably noticed a similar Doppler effect for sound waves. When a police car is approaching, the sound waves from its siren have a shorter wavelength and higher frequency than if the siren were at rest, and hence you hear a higher pitch. After the police car passes you and is moving away, you hear a lower pitch from the siren because the sound waves have a longer wavelength and a lower frequency.

Suppose that $\lambda_0$ is the wavelength of a particular spectral line from a light source that is not moving. It is the wavelength that you might look up in a reference book or determine in a laboratory experiment for this spectral line. If the source is moving, this particular spectral line is shifted to a different wavelength $\lambda$. The size of the wavelength shift is usually written as $\Delta\lambda$, where $\Delta\lambda = \lambda - \lambda_0$. Thus, $\Delta\lambda$ is the difference between the wavelength listed in reference books and the wavelength that you actually observe in the spectrum of a star or galaxy.



Doppler proved that the wavelength shift ($\Delta\lambda$) is governed by the following simple equation:

**Doppler shift equation**

$$\frac{\Delta\lambda}{\lambda_0} = \frac{v}{c}$$

$\Delta\lambda$ = wavelength shift

$\lambda_0$ = wavelength if source is not moving

$v$ = velocity of the source measured along the line of sight

$c$ = speed of light = $3.0 \times 10^5$ km/s

CAUTION! The capital Greek letter $\Delta$ (delta) is commonly used to denote a change in the value of a quantity. Thus, $\Delta\lambda$ is the change in the wavelength $\lambda$ due to the Doppler effect. It is *not* equal to a quantity $\Delta$ multiplied by a second quantity $\lambda$!

### Interpreting the Doppler Effect

The velocity determined from the Doppler effect is called **radial velocity,** because $v$ is the component of the star's motion parallel to our line of sight, or along the "radius" drawn from Earth to the star. Of course, a sizable fraction of a star's motion may be perpendicular to our line of sight. The speed of this transverse movement across the sky does not affect wavelengths if the speed is small compared with $c$. Box 5-6 includes two examples of calculations with radial velocity using the Doppler formula.

CAUTION! The redshifts and blueshifts of stars visible to the naked eye, or even through a small telescope, are only a small fraction of a nanometer. These tiny wavelength changes are far too small to detect visually. (Astronomers were able to detect the tiny Doppler shifts of starlight only after they had developed highly sensitive equipment for measuring wavelengths. This was done around 1890, a half-century after Doppler's original proposal.) So, if you see a star with a red color, it means that the star really is red; it does *not* mean that it is moving rapidly away from us.

The Doppler effect is an important tool in astronomy because it uncovers basic information about the motions of planets, stars, and galaxies. For example, the rotation of the planet Venus was deduced from the Doppler shift of radar waves reflected from its surface. Small Doppler shifts in the spectrum of sunlight have shown that the entire Sun is vibrating like an immense gong. The back-and-forth Doppler shifting of the spectral lines of certain stars reveals that these stars are being orbited by unseen companions; from this astronomers have discovered planets around other stars and massive objects that may be black holes. Astronomers also use the Doppler effect along with Kepler's third law to measure the masses of galaxies. These are but a few examples of how Doppler's discovery has empowered astronomers in their quest to understand the universe.

In this chapter we have glimpsed how much can be learned by analyzing light from the heavens. To analyze this light, however, it is first necessary to collect as much of it as possible, because most light sources in space are very dim. Collecting the faint light from distant objects is a key purpose of telescopes. In the next chapter we will describe both how telescopes work and how they are used.

## Key Words

*Terms preceded by an asterisk (*) are discussed in the Boxes.*

absolute zero, p. 104
absorption line spectrum, p. 112
atom, p. 104
*atomic number, p. 118
Balmer line, p. 116
Balmer series, p. 116
blackbody, p. 106
blackbody curve, p. 106
blackbody radiation, p. 106
blueshift, p. 122
Bohr orbits, p. 117
compound, p. 112
continuous spectrum, p. 112
*degrees Celsius, p. 105
*degrees Fahrenheit, p. 105
Doppler effect, p. 122
electromagnetic radiation, p. 101
electromagnetic spectrum, p. 102
electromagnetism, p. 100
electron, p. 109
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excited state, p. 120
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infrared radiation, p. 101
ionization, p. 120
*isotope, p. 118
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radial velocity, p. 123
radio waves, p. 101
redshift, p. 122
*solar constant, p. 108
spectral analysis, p. 111
spectral line, p. 111
spectroscopy, p. 113
spectrum (*plural* spectra), p. 99
Stefan-Boltzmann law, p. 107
ultraviolet radiation, p. 102
visible light, p. 101
watt, p. 107
wavelength, p. 101
wavelength of maximum emission, p. 105
Wien's law, p. 107
X rays, p. 102

## Key Ideas

**The Nature of Light:** Light is electromagnetic radiation. It has wavelike properties described by its wavelength $\lambda$ and frequency $\nu$, and travels through empty space at the constant speed $c = 3.0 \times 10^8$ m/s $= 3.0 \times 10^5$ km/s.

**Blackbody Radiation:** A blackbody is a hypothetical object that is a perfect absorber of electromagnetic radiation at all wavelengths. Stars closely approximate the behavior of blackbodies, as do other hot, dense objects.

• The intensities of radiation emitted at various wavelengths by a blackbody at a given temperature are shown by a blackbody curve.

• Wien's law states that the dominant wavelength at which a blackbody emits electromagnetic radiation is inversely proportional to the Kelvin temperature of the object: $\lambda_{max}$ (in meters) = (0.0029 K m)/$T$.

• The Stefan-Boltzmann law states that a blackbody radiates electromagnetic waves with a total energy flux $F$ directly proportional to the fourth power of the Kelvin temperature $T$ of the object: $F = \sigma T^4$.

**Photons:** An explanation of blackbody curves led to the discovery that light has particlelike properties. The particles of light are called photons.

• Planck's law relates the energy $E$ of a photon to its frequency $\nu$ or wavelength $\lambda$: $E = h\nu = hc/\lambda$, where $h$ is Planck's constant.

**Kirchhoff's Laws:** Kirchhoff's three laws of spectral analysis describe conditions under which different kinds of spectra are produced.

• A hot, dense object such as a blackbody emits a continuous spectrum covering all wavelengths.

• A hot, transparent gas produces a spectrum that contains bright (emission) lines.

• A cool, transparent gas in front of a light source that itself has a continuous spectrum produces dark (absorption) lines in the continuous spectrum.

**Atomic Structure:** An atom has a small dense nucleus composed of protons and neutrons. The nucleus is surrounded by electrons that occupy only certain orbits or energy levels.

• When an electron jumps from one energy level to another, it emits or absorbs a photon of appropriate energy (and hence of a specific wavelength).

• The spectral lines of a particular element correspond to the various electron transitions between energy levels in atoms of that element.

• Bohr's model of the atom correctly predicts the wavelengths of hydrogen's spectral lines.

**The Doppler Shift:** The Doppler shift enables us to determine the radial velocity of a light source from the displacement of its spectral lines.

• The spectral lines of an approaching light source are shifted toward short wavelengths (a blueshift); the spectral lines of a receding light source are shifted toward long wavelengths (a redshift).

• The size of a wavelength shift is proportional to the radial velocity of the light source relative to the observer.

## Questions

### Review Questions

1. When Jupiter is undergoing retrograde motion as seen from Earth, would you expect the eclipses of Jupiter's moons to occur several minutes early, several minutes late, or neither? Explain your answer.
2. Approximately how many times around Earth could a beam of light travel in one second?
3. How long does it take light to travel from the Sun to Earth, a distance of $1.50 \times 10^8$ km?
4. TUTORIAL 5.1 How did Newton show that a prism breaks white light into its component colors, but does not add any color to the light?
5. For each of the following wavelengths, state whether it is in the radio, microwave, infrared, visible, ultraviolet, X-ray, or gamma-ray portion of the electromagnetic spectrum. Explain your reasoning. (a) 2.6 μm, (b) 34 m, (c) 0.54 nm, (d) 0.0032 nm, (e) 0.620 μm, (f) 310 nm, (g) 0.012 m
6. What is meant by the frequency of light? How is frequency related to wavelength?
7. A cellular phone is actually a radio transmitter and receiver. You receive an incoming call in the form of a radio wave of frequency 880.65 MHz. What is the wavelength (in meters) of this wave?
8. A light source emits infrared radiation at a wavelength of 1150 nm. What is the frequency of this radiation?
9. (a) What is a blackbody? (b) In what way is a blackbody black? (c) If a blackbody is black, how can it emit light? (d) If you were to shine a flashlight beam on a perfect blackbody, what would happen to the light?
10. Why do astronomers find it convenient to use the Kelvin temperature scale in their work rather than the Celsius or Fahrenheit scale?
11. Explain why astronomers are interested in blackbody radiation.
12. TUTORIAL 5.3 Using Wien's law and the Stefan-Boltzmann law, explain the color and intensity changes that are observed as the temperature of a hot, glowing object increases.
13. If you double the Kelvin temperature of a hot piece of steel, how much more energy will it radiate per second?
14. The bright star Bellatrix in the constellation Orion has a surface temperature of 21,500 K. What is its wavelength of maximum emission in nanometers? What color is this star?
15. The bright star Antares in the constellation Scorpius (the Scorpion) emits the greatest intensity of radiation at a wavelength of 853 nm. What is the surface temperature of Antares? What color is this star?
16. (a) Describe an experiment in which light behaves like a wave. (b) Describe an experiment in which light behaves like a particle.
17. How is the energy of a photon related to its wavelength? What kind of photons carry the most energy? What kind of photons carry the least energy?
18. To emit the same amount of light energy per second, which must emit more photons per second: a source of red light, or a source of blue light? Explain your answer.
19. Explain how we know that atoms have massive, compact nuclei.
20. (a) Describe the spectrum of hydrogen at visible wavelengths. (b) Explain how Bohr's model of the atom accounts for the Balmer lines.



21. Why do different elements display different patterns of lines in their spectra?
22. What is the Doppler effect? Why is it important to astronomers?
23. If you see a blue star, what does its color tell you about how the star is moving through space? Explain your answer.

### Advanced Questions

*Questions preceded by an asterisk (*) involve topics discussed in the Boxes.*

> Problem-solving tips and tools
>
> You can find formulas in Box 5-1 for converting between temperature scales. Box 5-2 discusses how a star's radius, luminosity, and surface temperature are related. Box 5-3 shows how to use Planck's law to calculate the energy of a photon. To learn how to do calculations using the Doppler effect, see Box 5-6.

24. Your normal body temperature is 98.6°F. What kind of radiation do you predominantly emit? At what wavelength (in nm) do you emit the most radiation?
25. What is the temperature of the Sun's surface in degrees Fahrenheit?
26. What wavelength of electromagnetic radiation is emitted with greatest intensity by this book? To what region of the electromagnetic spectrum does this wavelength correspond?
27. Black holes are objects whose gravity is so strong that not even an object moving at the speed of light can escape from their surface. Hence, black holes do not themselves emit light. But it is possible to detect radiation from material falling *toward* a black hole. Calculations suggest that as this matter falls, it is compressed and heated to temperatures around $10^6$ K. Calculate the wavelength of maximum emission for this temperature. In what part of the electromagnetic spectrum does this wavelength lie?

*28. Use the value of the solar constant given in Box 5-2 and the distance from Earth to the Sun to calculate the luminosity of the Sun.

*29. The star Alpha Lupi (the brightest in the constellation Lupus, the Wolf) has a surface temperature of 21,600 K. How much more energy is emitted each second from each square meter of the surface of Alpha Lupi than from each square meter of the Sun's surface?

*30. Jupiter's moon Io has an active volcano named Pele whose temperature can be as high as 320°C. (a) What is the wavelength of maximum emission for the volcano at this temperature? In what part of the electromagnetic spectrum is this? (b) The average temperature of Io's surface is −150°C. Compared with a square meter of surface at this temperature, how much more energy is emitted per second from each square meter of Pele's surface?

*31. The bright star Sirius in the constellation of Canis Major (the Large Dog) has a radius of 1.67 $R_\odot$ and a luminosity of 25 $L_\odot$. (a) Use this information to calculate the energy flux at the surface of Sirius. (b) Use your answer in part (a) to calculate the surface temperature of Sirius. How does your answer compare to the value given in Box 5-2?

32. In Figure 5-13 you can see two distinct dark lines at the boundary between the orange and yellow parts of the Sun's spectrum (in the center of the third colored band from the top of the figure). The wavelengths of these dark lines are 588.99 and 589.59 nm. What do you conclude from this about the chemical composition of the Sun's atmosphere? (*Hint:* See Section 5-6.)
33. Instruments on board balloons and spacecraft detect 511-keV photons coming from the direction of the center of our Galaxy. (The prefix k means *kilo*, or thousand, so 1 keV = $10^3$ eV.) What is the wavelength of these photons? To what part of the electromagnetic spectrum do these photons belong?
34. (a) Calculate the wavelength of $P_\delta$ (P-delta), the fourth wavelength in the Paschen series. (b) Draw a schematic diagram of the hydrogen atom and indicate the electron transition that gives rise to this spectral line. (c) In what part of the electromagnetic spectrum does this wavelength lie?
35. (a) Calculate the wavelength of $H_\eta$ (H-eta), the spectral line for an electron transition between the $n = 7$ and $n = 2$ orbits of hydrogen. (b) In what part of the electromagnetic spectrum does this wavelength lie? Use this to explain why Figure 5-21 is labeled R I V U X G.
36. Certain interstellar clouds contain a very cold, very thin gas of hydrogen atoms. Ultraviolet radiation with any wavelength shorter than 91.2 nm cannot pass through this gas; instead, it is absorbed. Explain why.
37. (a) Can a hydrogen atom in the ground state absorb an H-alpha ($H_\alpha$) photon? Explain why or why not. (b) Can a hydrogen atom in the $n = 2$ state absorb a Lyman-alpha ($L_\alpha$) photon? Explain why or why not.
38. An imaginary atom has just 3 energy levels: 0 eV, 1 eV, and 3 eV. Draw an energy-level diagram for this atom. Show all possible transitions between these energy levels. For each transition, determine the photon energy and the photon wavelength. Which transitions involve the emission or absorption of visible light?
39. The star cluster NGC 346 and nebula shown in Figure 5-18 are located within the Small Magellanic Cloud (SMC), a small galaxy that orbits our Milky Way Galaxy. The SMC and the stars and gas within it are moving away from us at 158 km/s. At what wavelength does the red $H_\alpha$ line of hydrogen (which causes the color of the nebula) appear in the nebula's spectrum?
40. The wavelength of $H_\beta$ in the spectrum of the star Megrez in the Big Dipper (part of the constellation Ursa Major, the Great Bear) is 486.112 nm. Laboratory measurements demonstrate that the normal wavelength of this spectral line is 486.133 nm. Is the star coming toward us or moving away from us? At what speed?
41. You are given a traffic ticket for going through a red light (wavelength 700 nm). You tell the police officer that because you were approaching the light, the Doppler effect caused a blueshift that made the light appear green (wavelength 500 nm). How fast would you have had to be going for this to be true? Would the speeding ticket be justified? Explain your answer.

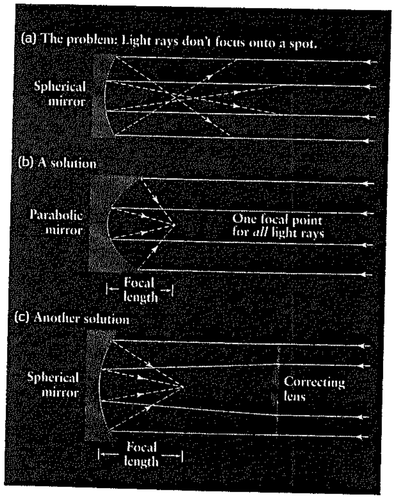
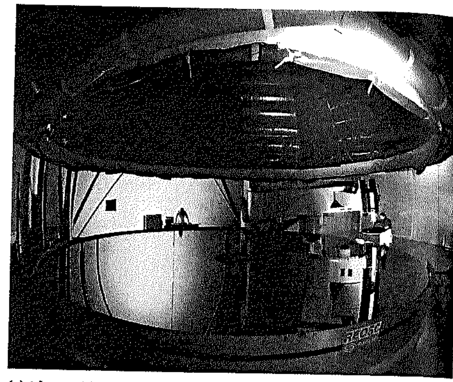
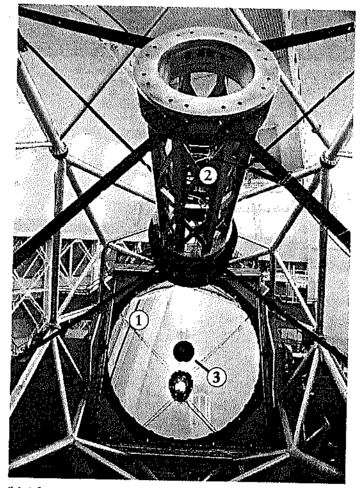



Figure 6-13

Spherical Aberration **(a)** Different parts of a spherically concave mirror reflect light to slightly different points. This effect, called spherical aberration, causes image blurring. This difficulty can be corrected by either **(b)** using a parabolic mirror or **(c)** using a correcting lens in front of the mirror.



(a) A large objective mirror




(b) A large Cassegrain telescope

Figure 6-14 R I V U X G

Reflecting Telescopes **(a)** This photograph shows technicians preparing an objective mirror 8.2 meters in diameter for the European Southern Observatory in Chile. The mirror was ground to a curved shape with a remarkable precision of 8.5 nanometers. **(b)** This view of the Gemini North telescope shows its 8.1-meter objective mirror (1). Light incident on this mirror is reflected toward the 1.0-meter secondary mirror (2), then through the hole in the objective mirror (3) to the Cassegrain focus (see Figure 6-11*b*). (a: SAGEM; b: NOAO/AURA/NSF)

from a defect called **coma,** wherein star images far from the center of the field of view are elongated to look like tiny teardrops. A different approach is to use a spherical mirror, thus minimizing coma, and to place a thin correcting lens at the front of the telescope to eliminate spherical aberration (Figure 6-13*c*). This approach is only used on relatively small reflecting telescopes for amateur astronomers.

## The Largest Reflectors

WEB LINK 6.2 There are over a dozen optical reflectors in operation with primary mirrors between 8 meters (26.2 feet) and 11 meters (36.1 feet) in diameter. Figure 6-14*a* shows the objective mirror of one of the four Very Large Telescope (VLT) units in Chile, and Figure 6-14*b* shows the objective and secondary mirrors of the Gemini North telescope in Hawaii. A near-twin of Gemini North, called Gemini South, is in Cerro Pachón, Chile. These twins allow astronomers to observe both the northern and southern parts of the celestial sphere with essentially the same state-of-the-art instrument. Two other "twins" are the side-by-side 8.4-m objective mirrors of the Large Binocular Telescope in Arizona. Combining the light from these two mirrors gives double the light-gathering power, equivalent to a single 11.8-m mirror.



Several other reflectors around the world have objective mirrors between 3 and 6 meters in diameter, and dozens of smaller but still powerful telescopes have mirrors in the range of 1 to 3 meters. There are thousands of professional astronomers, each of whom has several ongoing research projects, and thus the demand for all of these telescopes is high. On any night of the year, nearly every research telescope in the world is being used to explore the universe.

## 6-3 Telescope images are degraded by the blurring effects of the atmosphere and by light pollution

TUTORIAL 6.1 In addition to providing a brighter image, a large telescope also helps achieve a second major goal: It produces star images that are sharp and crisp. A quantity called **angular resolution** gauges how well fine details can be seen. Poor angular resolution causes star images to be fuzzy and blurred together.

To determine the angular resolution of a telescope, pick out two adjacent stars whose separate images are just barely discernible (Figure 6-15). The angle θ (the Greek letter theta) between these stars is the telescope's angular resolution; the *smaller* that angle, the finer the details that can be seen and the sharper the image.

When you are asked to read the letters on an eye chart, what's being measured is the angular resolution of your eye. If you have 20/20 vision, the angular resolution θ of your eye is about 1 arcminute, or 60 arcseconds. (You may want to review the definitions of these angular measures in Section 1-5.) Hence, with the naked eye it is impossible to distinguish two stars less than 1 arcminute apart or to see details on the Moon with an angular size smaller than this. All the planets have angular sizes (as seen from Earth) of 1 arcminute or less, which is why they appear as featureless points of light to the naked eye.

### Limits to Angular Resolution

One factor limiting angular resolution is **diffraction,** which is the tendency of light waves to spread out when they are confined to a small area like the lens or mirror of a telescope. (A rough analogy is the way water exiting a garden hose sprays out in a wider angle when you cover part of the end of the hose with your thumb.) As a result of diffraction, a narrow beam of light tends to spread out within a telescope's optics, thus blurring the image. If diffraction were the only limit, the angular resolution of a telescope would be given by the formula

**Diffraction-limited angular resolution**

$$\theta = 2.5 \times 10^5 \frac{\lambda}{D}$$

θ = diffraction-limited angular resolution of a telescope, in arcseconds

λ = wavelength of light, in meters

$D$ = diameter of telescope objective, in meters

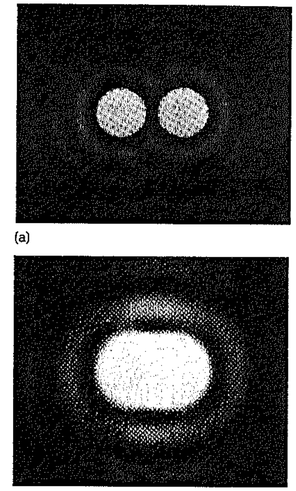



Figure 6-15 R I V U X G

Angular Resolution The angular resolution of a telescope indicates the sharpness of the telescope's images. **(a)** This telescope view shows two sources of light whose angular separation is greater than the angular resolution. **(b)** The light sources have been moved together so that their angular separation is equal to the angular resolution. If the sources were moved any closer together, the telescope image would show them as a single source.

For a given wavelength of light, using a telescope with an objective of *larger* diameter $D$ *reduces* the amount of diffraction and makes the angular resolution θ *smaller* (and hence better). For example, with red light with wavelength 640 nm, or $6.4 \times 10^{-7}$ m, the diffraction-limited resolution of an 8-meter telescope (see Figure 6-15) would be

$$\theta = (2.5 \times 10^5)\,\frac{6.4 \times 10^{-7}\ \text{m}}{8\ \text{m}} = 0.02\ \text{arcsec}$$

In practice, however, ordinary optical telescopes cannot achieve such fine angular resolution. The problem is that turbulence in the air causes star images to jiggle around and twinkle. Even through the largest telescopes, a star still looks like a tiny blob rather than a pinpoint of light. A measure of the limit that atmospheric turbulence places on a telescope's resolution is called the **seeing disk.** This disk is the angular diameter of a star's image broadened by turbulence. The size of the seeing disk varies from one observatory site to another and from one night to another. At the observatories on Kitt Peak in Arizona and Cerro Tololo in Chile, the seeing disk is typically around 1 arcsec. Some

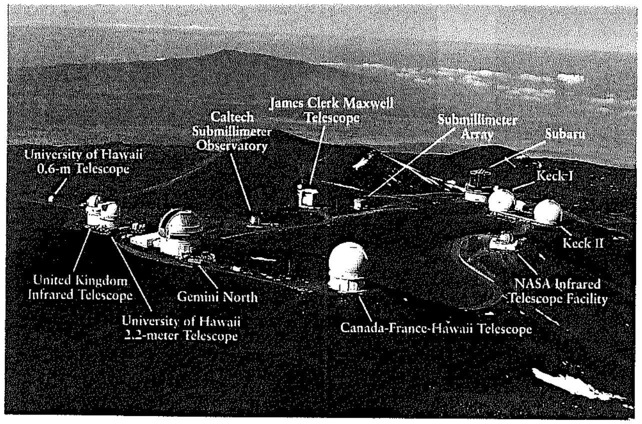



**Figure 6-16** R I V U X G

The Telescopes of Mauna Kea The summit of Mauna Kea—an extinct Hawaiian volcano that reaches more than 4100 m (13,400 ft) above the waters of the Pacific—has nighttime skies that are unusually clear, still, and dark. To take advantage of these superb viewing conditions, Mauna Kea has become the home of many powerful telescopes. (Richard J. Wainscoat, University of Hawaii)

of the very best conditions in the world can be found at the observatories atop Mauna Kea in Hawaii, where the seeing disk is often as small as 0.5 arcsec. These great conditions are one reason why so many telescopes have been built there (Figure 6-16).

## Active Optics and Adaptive Optics

In many cases the angular resolution of a telescope is even worse than the limit imposed by the seeing disk. This occurs if the objective mirror deforms even slightly due to variations in air temperature or flexing of the telescope mount. To combat this, many large telescopes are equipped with an **active optics** system. Such a system adjusts the mirror shape every few seconds to help keep the telescope in optimum focus and properly aimed at its target.

Changing the mirror shape is also at the heart of a more refined technique called **adaptive optics**. The goal of this technique is to compensate for atmospheric turbulence, so that the angular resolution can be smaller than the size of the seeing disk and can even approach the theoretical limit set by diffraction. Turbulence causes the image of a star to "dance" around erratically. In an adaptive optics system, sensors monitor this dancing motion 10 to 100 times per second, and a powerful computer rapidly calculates the mirror shape needed to compensate. Fast-acting mechanical devices called *actuators* then deform the mirror accordingly, at a much faster rate than in an active optics system. In some adaptive optics systems, the actuators deform a small secondary mirror rather than the large objective mirror.

**Adaptive optics produces sharper images by "undoing" atmospheric turbulence**

One difficulty with adaptive optics is that a fairly bright star must be in or near the field of the telescope's view to serve as a "target" for the sensors that track atmospheric turbulence. This is seldom the case, since the field of view of most telescopes is rather narrow. Astronomers get around this limitation by shining a laser beam toward a spot in the sky near the object to be observed (Figure 6-17). The laser beam causes atoms in the upper atmosphere to glow, making an artificial "star." The light that comes down to Earth from this "star" travels through the same part of our atmosphere as the light from the object being observed, so its image in the telescope will "dance" around in the same erratic way as the image of a real star.

Figure 6-18 shows the dramatic improvement in angular resolution possible with adaptive optics. Images made with adaptive optics are nearly as sharp as if the telescope were in the vacuum of space, where there is no atmospheric distortion whatsoever and the only limit on angular resolution is diffraction. A number of large telescopes are now being used with adaptive optics systems.

**CAUTION!** The images in Figure 6-18 are **false color** images: They do not represent the true color of the stars shown. False color is often used when the image is made using wavelengths that the eye cannot detect, as with the infrared images in Figure 6-18. A different use of false color is to indicate the relative brightness of different parts of the image, as in the infrared image of a person in Figure 5-10. Throughout this book, we'll always point out when false color is used in an image.

## Interferometry

Several large observatories are developing a technique called **interferometry** that promises to further improve the angular resolution of telescopes. The idea is to have two widely separated telescopes observe the same object simultaneously, then use fiber optic cables to "pipe" the light signals from each telescope to a central location where they "interfere" or blend together. This method makes the combined signal sharp and clear. The effective resolution of such a combination of telescopes is equivalent to that of one giant telescope with a diameter equal to the **baseline**, or distance between the two telescopes. For example, the Keck I



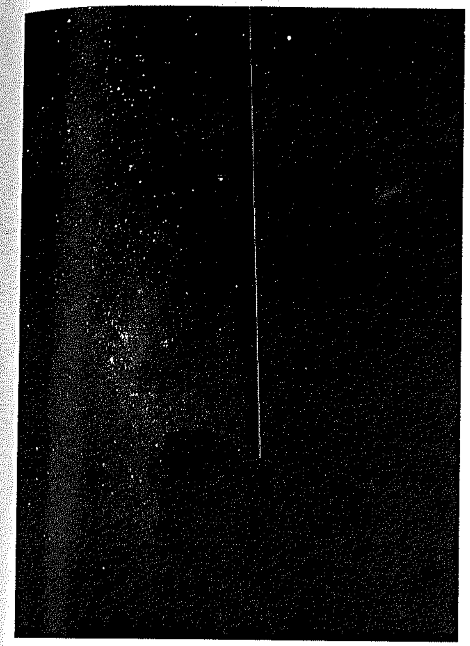


**Figure 6-17** R I V U X G

Creating an Artificial "Star" A laser beam shines upward from Yepun, an 8.2-meter telescope at the European Southern Observatory in the Atacama Desert of Chile. (Figure 6-14*a* shows the objective mirror for this telescope.) The beam strikes sodium atoms that lie about 90 km (56 miles) above Earth's surface, causing them to glow and make an artificial "star." Tracking the twinkling of this "star" makes it possible to undo the effects of atmospheric turbulence on telescope images. (European Southern Observatory)

and Keck II telescopes atop Mauna Kea (Figure 6-16) are 85 meters apart, so when used as an interferometer the angular resolution is the same as a single 85-meter telescope.

Interferometry has been used for many years with radio telescopes (which we will discuss in Section 6-6), but is still under development with telescopes for visible light or infrared wavelengths. Astronomers are devoting a great deal of effort to this development because the potential rewards are great. For example, the Keck I and II telescopes used together should give an angular resolution as small as 0.005 arcsec, which corresponds to being able to read the bottom row on an eye chart 36 km (22 miles) away!

## Light Pollution

Light from city street lamps and from buildings also degrades telescope images. This **light pollution** illuminates the sky, making it more difficult to see the stars. You can appreciate the problem if you have ever looked at the night sky from a major city. Only a few of the very brightest stars can be seen, as against the thousands that can be seen with the naked eye in the desert or the mountains. To avoid light pollution, observatories are built in remote locations far from any city lights.

Unfortunately, the expansion of cities has brought light pollution to observatories that in former times had none. As an example, the growth of Tucson, Arizona, has had deleterious effects on observations at the nearby Kitt Peak National Observatory. Efforts have been made to have cities adopt light fixtures that provide safe illumination for their citizens but produce little light pollution. These efforts have met with only mixed success.

One factor over which astronomers have absolutely no control is the weather. Optical telescopes cannot see through clouds, so it is important to build observatories where the weather is usually good. One advantage of mountaintop observatories such as Mauna Kea is that most clouds form at altitudes below the observatory, giving astronomers a better chance of having clear skies.

In many ways the best location for a telescope is in orbit around Earth, where it is unaffected by weather, light pollution, or atmospheric turbulence. We will discuss orbiting telescopes in Section 6-7.

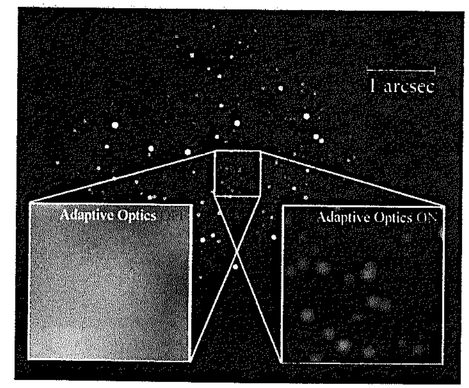



**Figure 6-18** R I V U X G

Using Adaptive Optics to "Unblur" Telescope Images The two false-color, inset images show the same 1-arcsecond-wide region of the sky at infrared wavelengths as observed with the 10.0-m Keck II telescope on Mauna Kea (see Figure 6-16). Without adaptive optics, it is impossible to distinguish individual stars in this region. With adaptive optics turned on, more than two dozen stars can be distinguished. (UCLA Galactic Center Group)

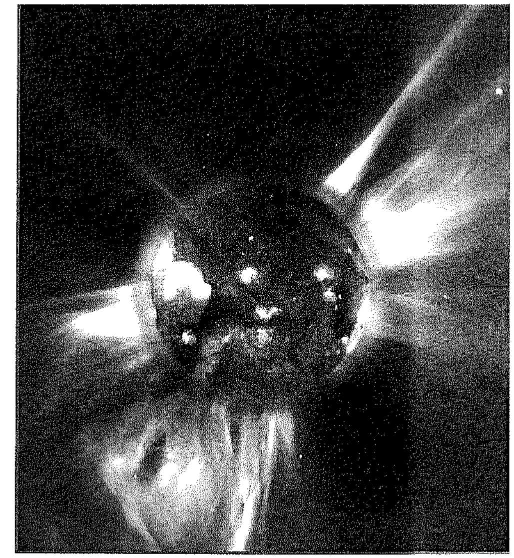
## Pluto and the Kuiper Belt *(continued)*

Mercury, Venus, Earth and Mars are terrestrial planets whose compositions are dominated by rock. Jupiter and Saturn are gas giant planets dominated by their hydrogen and helium envelopes. Uranus and Neptune are ice giant planets dominated by gases other than hydrogen and helium. The trans-Neptunian objects or "ice dwarf planets," as some are calling them, are probably composed of large amounts of volatiles such as methane ice and water ice.

The International Astronomical Union (IAU) recently choose to use a combination of the second and third definitions above to define a planet. That is, a planet is an object that is both spherical and has a unique orbit in which it is gravitationally dominant. An object that only satisfies definition two above and not three is now being called a dwarf planet, of which Ceres, Pluto and Eris qualify, as well as several more objects in the Kuiper belt. In reality, it's the public that must accept this definition, and we will only know if that is the case a few generations from now.

Defining the word "planet" is like defining what an ocean is. At first it appears to be a simple term, but is very hard to precisely define. Further planet discoveries will be made and there are sure to be borderline cases. This is just the way nature is; it does not have little bins to nicely classify objects, but there is usually a continuous array of objects. The important thing to take away from all of this is to understand this array of objects by using our rapidly expanding scientific knowledge and understanding our place in the solar system.

# 16 Our Star, the Sun



R I V U X G

A composite view of the Sun (at ultraviolet wavelengths) and an upheaval in the Sun's outer atmosphere, or corona (at visible wavelengths). (SOHO/LASCO/EIT/ESA/NASA)

The Sun is by far the brightest object in the sky. By earthly standards, the temperature of its glowing surface is remarkably high, about 5800 K. Yet there are regions of the Sun that reach far higher temperatures of tens of thousands or even millions of kelvins. Gases at such temperatures emit ultraviolet light, which makes them prominent in the accompanying image from an ultraviolet telescope in space. Some of the hottest and most energetic regions on the Sun spawn immense disturbances and can propel solar material across space to reach Earth and other planets.

In recent decades, we have learned that the Sun shines because at its core hundreds of millions of tons of hydrogen are converted to helium every second. We have confirmed this picture by detecting the by-products of this transmutation—strange, ethereal particles called neutrinos—streaming outward from the Sun into space. We have discovered that the Sun has a surprisingly violent atmosphere, with a host of features such as sunspots whose numbers rise and fall on a predictable 11-year cycle. By studying the Sun's vibrations, we have begun to probe beneath its surface into hitherto unexplored realms. And we have just begun to investigate how changes in the Sun's activity can affect Earth's environment as well as our technological society.

## Learning Goals

*By reading the sections of this chapter, you will learn*

16-1 The source of the Sun's heat and light

16-2 How scientists model the Sun's internal structure

16-3 How the Sun's vibrations reveal what lies beneath its glowing surface

16-4 How scientists are able to probe the Sun's energy-generating core

16-5 Why the gaseous Sun appears to have a sharp outer edge

16-6 Why the upper regions of the solar atmosphere have an emission spectrum

16-7 The relationship between the Sun's corona and the solar wind

16-8 The nature of sunspots

16-9 The connection between sunspots and the Sun's magnetic field

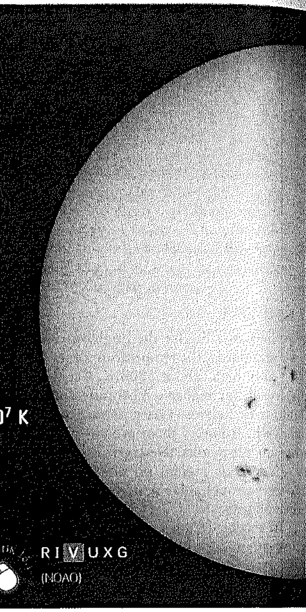
16-10 How magnetic reconnection can power immense solar eruptions

## 16-1 The Sun's energy is generated by thermonuclear reactions in its core

The Sun is the largest member of the solar system. It has almost a thousand times more mass than all the planets, moons, asteroids, comets, and meteoroids put together. But the Sun is also a star. In fact, it is a remarkably typical star, with a mass, size, surface temperature, and chemical composition that are roughly midway between the extremes exhibited by the myriad other stars in the heavens. 16-1 lists essential data about the Sun.

## Table 16-1 Sun Data

| | |
|---|---|
| Distance from Earth: | Mean: 1 AU = 149,598,000 km |
| | Maximum: 152,000,000 km |
| | Minimum: 147,000,000 km |
| Light travel time to Earth: | 8.32 min |
| Mean angular diameter: | 32 arcmin |
| Radius: | 696,000 km = 109 Earth radii |
| Mass: | $1.9891 \times 10^{30}$ kg = $3.33 \times 10^5$ Earth masses |
| Composition (by mass): | 74% hydrogen, 25% helium, 1% other elements |
| Composition (by number of atoms): | 92.1% hydrogen, 7.8% helium, 0.1% other elements |
| Mean density: | 1410 kg/m$^3$ |
| Mean temperatures: | Surface: 5800 K; Center: $1.55 \times 10^7$ K |
| Luminosity: | $3.90 \times 10^{26}$ W |
| Distance from center of Galaxy: | 8000 pc = 26,000 ly |
| Orbital period around center of Galaxy: | 220 million years |
| Orbital speed around center of Galaxy: | 220 km/s |




### Solar Energy

For most people, what matters most about the Sun is the energy that it radiates into space. Without the Sun's warming rays, our atmosphere and oceans would freeze into an icy layer coating a desperately cold planet, and life on Earth would be impossible. To understand why we are here, we must understand the nature of the Sun.

Why is the Sun such an important source of energy? An important part of the answer is that the Sun has a far higher surface temperature than any of the planets or moons. The Sun's spectrum is close to that of an idealized blackbody with a temperature of 5800 K (see Sections 5-3 and 5-4, especially Figure 5-12). Thanks to this high temperature, each square meter of the Sun's surface emits a tremendous amount of radiation, principally at visible wavelengths. Indeed, the Sun is the only object in the solar system that emits substantial amounts of visible light. The light that we see from the Moon and planets is actually sunlight that struck those worlds and was reflected toward Earth.

The Sun's size also helps us explain its tremendous energy output. Because the Sun is so large, the total number of square meters of radiating surface—that is, its surface area—is immense. Hence, the total amount of energy emitted by the Sun each second, called its **luminosity**, is very large indeed or about $3.9 \times 10^{26}$ watts, ($3.9 \times 10^{26}$ joules of energy emitted per second.) (We discussed the relation among the Sun's surface temperature, radius, and luminosity in Box 5-2.) Astronomers denote the Sun's luminosity by the symbol $L_{\odot}$. A circle with a dot in the center is the astronomical symbol for the Sun and was also used by ancient astrologers.

### The Source of the Sun's Energy: Early Ideas

These ideas lead us to a more fundamental question: What keeps the Sun's visible surface so hot? Or, put another way, what is the fundamental source of the tremendous energies that the Sun radiates into space? For centuries, this question was one of the greatest mysteries in science. The mystery deepened in the nineteenth century, when geologists and biologists found convincing evidence that life has existed on Earth for at least several hundred million years. (We now know that Earth is 4.56 billion years old and that life had existed on it for most of its history.) Since life as we know it depends crucially on sunlight, the Sun must be as old. The source of the Sun's energy posed a severe problem for physicists. What source of energy could have kept the Sun shining for so long (Figure 16-1)?

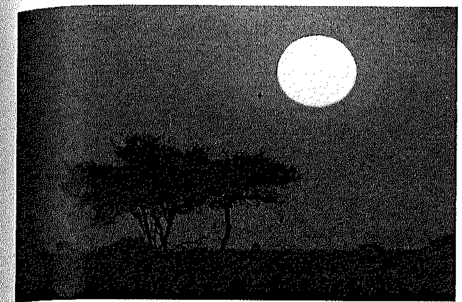


Figure 16-1 RIVUXG
The Sun The Sun's visible surface has a temperature of about 5800 K. At this temperature, all solids and liquids vaporize to form gases. It was only in the twentieth century that scientists discovered what has kept the Sun so hot for billions of years: the thermonuclear fusion of hydrogen nuclei in the Sun's core. (Jeremy Woodhouse/PhotoDisc)

One attempt to explain solar energy was made in the mid-1800s by the English physicist Lord Kelvin (for whom the temperature scale is named) and the German scientist Hermann von Helmholtz. They argued that the tremendous weight of the Sun's outer layers should cause the Sun to contract gradually, compressing its interior gases. Whenever a gas is compressed, its temperature rises. (You can demonstrate this with a bicycle pump: As you pump air into a tire, the temperature of the air increases and the pump becomes warm to the touch.) Kelvin and Helmholtz thus suggested that gravitational contraction could cause the Sun's gases to become hot enough to radiate energy out into space.

This process, called *Kelvin-Helmholtz contraction,* actually does occur during the earliest stages of the birth of a star like the Sun (see Section 8-4). But Kelvin-Helmholtz contraction cannot be the major source of the Sun's energy today. If it were, the Sun would have had to be much larger in the relatively recent past. Helmholtz's own calculations showed that the Sun could have started its initial collapse from the solar nebula no more than about 25 million years ago. But the geological and fossil record shows that Earth is far older than that, and so the Sun must be as well. Hence, this model of a Sun that shines because it shrinks cannot be correct.

On Earth, a common way to produce heat and light is by burning fuel, such as a log in a fireplace or coal in a power plant. Is it possible that a similar process explains the energy released by the Sun? The answer is no, because this process could not continue for a long enough time to explain the age of Earth. The chemical reactions involved in burning release roughly $10^{-19}$ joule of energy per atom. Therefore, the number of atoms that would have to undergo chemical reactions each second to generate the Sun's luminosity of $3.9 \times 10^{26}$ joules per second is approximately

$$\frac{3.9 \times 10^{26} \text{ joules per second}}{10^{-19} \text{ joule per atom}} = 3.9 \times 10^{45} \text{ atoms per second}$$

From its mass and chemical composition, we know that the Sun contains about $10^{57}$ atoms. Thus, the length of time that would be required to consume the entire Sun by burning is

$$\frac{10^{57} \text{ atoms}}{3.9 \times 10^{45} \text{ atoms per second}} = 3 \times 10^{11} \text{ seconds}$$

There are about $3 \times 10^7$ seconds in a year. Hence, in this model, the Sun would burn itself out in a mere 10,000 ($10^4$) years! This period of time is far shorter than the known age of Earth, so chemical reactions also cannot explain how the Sun shines.

### The Source of the Sun's Energy: Discovering Thermonuclear Fusion

The source of the Sun's luminosity could be explained if there were a process that was like burning but released much more energy per atom. Then the rate at which atoms would have to be consumed would be far less, and the lifetime of the Sun could be long enough to be consistent with the known age of Earth. Albert Einstein discovered the key to such a process in 1905. According to his *special theory of relativity,* a quantity $m$ of mass can in principle be converted into an amount of energy $E$ according to a now-famous equation:

**Einstein's mass-energy equation**

$$E = mc^2$$

$m$ = quantity of mass, in kg

$c$ = speed of light = $3 \times 10^8$ m/s

$E$ = amount of energy into which the mass can be converted, in joules

The speed of light $c$ is a large number, so $c^2$ is huge. Therefore, a small amount of matter can release an awesome amount of energy.

Inspired by Einstein's ideas, astronomers began to wonder if the Sun's energy output might come from the conversion of matter into energy. The Sun's low density of 1410 kg/m$^3$ indicates that it must be made of the very lightest atoms, primarily hydrogen and helium. In the 1920s, the British astronomer Arthur Eddington showed that temperatures near the center of the Sun must be so high that atoms become completely ionized. Hence, at the Sun's center we expect to find hydrogen nuclei and electrons flying around independent of each other.

**Ideas from relativity and nuclear physics led to an understanding of how the Sun shines**

Another British astronomer, Robert Atkinson, suggested that under these conditions hydrogen nuclei could fuse together to produce helium nuclei in a *nuclear reaction* that transforms a tiny amount of mass

**BOX 16-1** Tools of the Astronomer's Trade

## Converting Mass into Energy

The *Cosmic Connections* figure shows the steps involved in the thermonuclear fusion of hydrogen at the Sun's center. In these steps, four protons are converted into a single nucleus of $^4$He, an isotope of helium with two protons and two neutrons. (As we saw in Box 5-5, different isotopes of the same element have the same number of protons but different numbers of neutrons.) The reaction depicted in the *Cosmic Connections: The Proton-Proton Chain* Step 1 also produces a neutral, nearly massless particle called the *neutrino*. Neutrinos respond hardly at all to ordinary matter, so they travel almost unimpeded through the Sun's massive bulk. Hence, the energy that neutrinos carry is quickly lost into space. This loss is not great, however, because the neutrinos carry relatively little energy. (See Section 16-4 for more about these curious particles.)

Most of the energy released by thermonuclear fusion appears in the form of gamma-ray photons. The energy of these photons remains trapped within the Sun for a long time, thus maintaining the Sun's intense internal heat. Some gamma-ray photons are produced by the reaction shown as Step 2 in the *Cosmic Connections* figure. Others appear when an electron in the Sun's interior annihilates a positively charged electron, or **positron**, which is a by-product of the reaction shown in Step 1 in the *Cosmic Connections* figure. An electron and a positron are respectively matter and antimatter, and they convert entirely into energy when they meet. (You may have thought that "antimatter" was pure science fiction. In fact, tremendous amounts of antimatter are being created and annihilated in the Sun as you read these words.)

We can summarize the thermonuclear fusion of hydrogen as follows:

$$4\,^1\text{H} \rightarrow {}^4\text{He} + \text{neutrinos} + \text{gamma-ray photons}$$

To calculate how much energy is released in this process, we use Einstein's mass-energy formula: The energy released is equal to the amount of mass consumed multiplied by $c^2$, where $c$ is the speed of light. To see how much mass is consumed, we compare the combined mass of four hydrogen atoms (the ingredients) to the mass of one helium atom (the product):

$$\begin{aligned} 4 \text{ hydrogen atoms} &= 6.693 \times 10^{-27} \text{ kg} \\ -1 \text{ helium atom} &= 6.645 \times 10^{-27} \text{ kg} \\ \hline \text{Mass lost} &= 0.048 \times 10^{-27} \text{ kg} \end{aligned}$$

Thus, a small fraction (0.7%) of the mass of the hydrogen going into the nuclear reaction does not show up in the mass of the helium. This lost mass is converted into an amount of energy $E = mc^2$:

$$\begin{aligned} E = mc^2 &= (0.048 \times 10^{-27} \text{ kg})(3 \times 10^8 \text{ m/s})^2 \\ &= 4.3 \times 10^{-12} \text{ joule} \end{aligned}$$

This amount of energy is released by the formation of a single helium atom. It would light a 10-watt lightbulb for almost one-half of a trillionth of a second.

EXAMPLE: How much energy is released when 1 kg of hydrogen is converted to helium?

Situation: We are given the initial mass of hydrogen. We know that a fraction of the mass is lost when the hydrogen undergoes fusion to make helium; our goal is to find the quantity of energy into which this lost mass is transformed.

Tools: We use the equation $E = mc^2$ and the result that 0.7% of the mass is lost when hydrogen is converted into hydrogen.

Answer: When 1 kilogram of hydrogen is converted to helium, the amount of mass lost is 0.7% of 1 kg, or 0.007 kg. (This means that 0.993 kg of helium is produced.) Using Einstein's equation, we find that this missing 0.007 kg of matter is transformed into an amount of energy equal to

$$E = mc^2 = (0.007 \text{ kg})(3 \times 10^8 \text{ m/s})^2 = 6.3 \times 10^{14} \text{ joules}$$

Review: The energy released by the fusion of 1 kilogram of hydrogen is the same as that released by burning *20,000 metric tons* ($2 \times 10^7$ kg) of coal! Hydrogen fusion is a *much* more efficient energy source than ordinary burning.

The Sun's luminosity is $3.9 \times 10^{26}$ joules per second. To generate this much power, hydrogen must be consumed at a rate of

$$\frac{3.9 \times 10^{26} \text{ joules per second}}{6.3 \times 10^{14} \text{ joules per kilogram}} = 6 \times 10^{11} \text{ kilograms per second}$$

That is, the Sun converts 600 million metric tons of hydrogen into helium every second.

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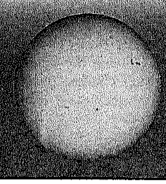
into a large amount of energy. Experiments in the laboratory using individual nuclei show that such reactions can indeed take place. The process of converting hydrogen into helium is called **hydrogen fusion**. (It is also sometimes called *hydrogen burning*, even though nothing is actually burned in the conventional sense. Ordinary burning involves chemical reactions that rearrange the outer electrons of atoms but have no effect on the atoms' nuclei.) Hydrogen fusion provides the devastating energy released in a hydrogen bomb (see Figure 1-6).

The fusing together of nuclei is also called **thermonuclear fusion**, because it can take place only at extremely high temperatures. The reason is that all nuclei have a positive electric charge

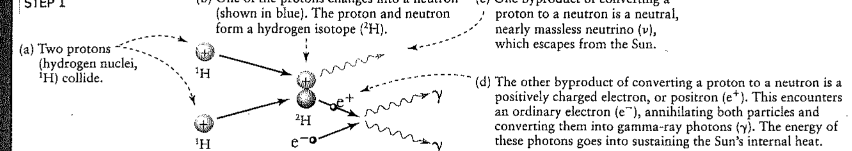
# COSMIC CONNECTIONS The Proton-Proton Chain

The most common form of hydrogen fusion in the Sun involves three steps, each of which releases energy.

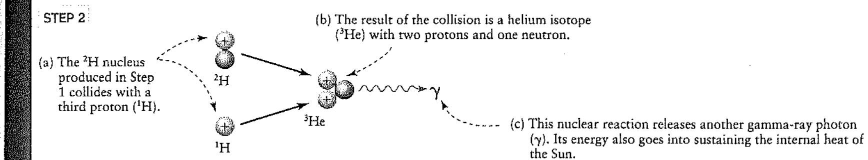
Hydrogen fusion in the Sun usually takes place in a sequence of steps called the proton-proton chain. Each of these steps releases energy that heats the Sun and gives it its luminosity.



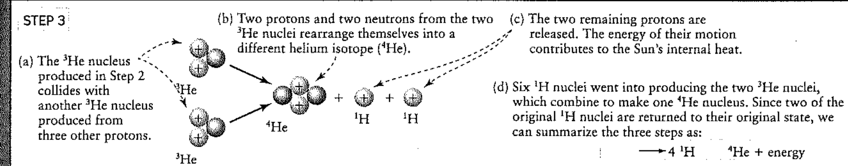
**STEP 1**



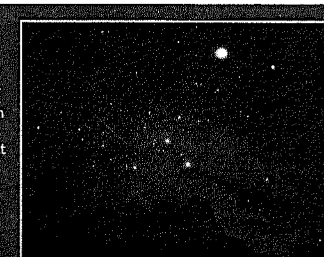

**STEP 2**




**STEP 3**




Hydrogen fusion also takes place in all of the stars visible to the naked eye. (Fusion follows a different sequence of steps in the most massive stars, but the net result is the same.)



and so tend to repel one another. But in the extreme heat and pressure at the Sun's center, hydrogen nuclei (protons) are moving so fast that they can overcome their electric repulsion and actually touch one another. When that happens, thermonuclear fusion can take place.

ANALOGY You can think of protons as tiny electrically charged spheres that are coated with a very powerful glue. If the spheres are not touching, the repulsion between their charges pushes them apart. But if the spheres are forced into contact, the strength of the glue "fuses" them together.

CAUTION! Be careful not to confuse thermonuclear fusion with the similar-sounding process of *nuclear fission*. In nuclear fusion, energy is released by joining or fusing together nuclei of lightweight atoms such as hydrogen. In nuclear fission, by contrast, the nuclei of very massive atoms such as uranium or plutonium release energy by fragmenting into smaller nuclei. Nuclear power plants produce energy using fission, not fusion. (Generating power using fusion has been a goal of researchers for decades, but no one has yet devised a commercially viable way to do this.)

### Converting Hydrogen to Helium

We learned in Section 5-8 that the nucleus of a hydrogen atom (H) consists of a single proton. The nucleus of a helium atom (He) consists of two protons and two neutrons. In the nuclear process that Atkinson described, four hydrogen nuclei combine to form one helium nucleus, with a concurrent release of energy:

$$4\ \text{H} \rightarrow \text{He} + \text{energy}$$

In several separate reactions, two of the four protons are changed into neutrons, and eventually combine with the remaining protons to produce a helium nucleus. This sequence of reactions is called the **proton-proton chain.** The *Cosmic Connections* figure depicts the proton-proton chain in detail.

Each time this process takes place, a small fraction (0.7%) of the initial combined mass of the hydrogen nuclei does not show up in the final mass of the helium nucleus. This "lost" mass is converted into energy. Box 16-1 describes how to use Einstein's mass-energy equation to calculate the amount of energy released.

CAUTION! You may have heard the idea that mass is always conserved (that is, it is neither created nor destroyed), or that energy is always conserved in a reaction. Einstein's ideas show that neither of these statements is quite correct, because mass can be converted into energy and vice versa. A more accurate statement is that the total amount of mass *plus* energy is conserved. Hence, the destruction of mass in the Sun does not violate any laws of nature.

For every four hydrogen nuclei converted into a helium nucleus, $4.3 \times 10^{-12}$ joule of energy is released. This amount of energy may seem tiny, but it is about $10^7$ times larger than the amount of energy released in a typical chemical reaction, such as occurs in ordinary burning. Thus, thermonuclear fusion explains how the Sun could have been shining for billions of years.

To produce the Sun's luminosity of $3.9 \times 10^{26}$ joules per second, $6 \times 10^{11}$ kg (600 million metric tons) of hydrogen must be converted into helium each second. This rate is prodigious, but there is a literally astronomical amount of hydrogen in the Sun. In particular, the Sun's core contains enough hydrogen to have been giving off energy at the present rate for as long as the solar system has existed, about 4.56 billion years, and to continue doing so for more than 6 billion years into the future.

LOOKING DEEPER 16.1 The proton-proton chain is also the energy source for many of the stars in the sky. In stars with central temperatures that are much hotter than that of the Sun, however, hydrogen fusion proceeds according to a different set of nuclear reactions, called the **CNO cycle**, in which carbon, nitrogen, and oxygen nuclei absorb protons to produce helium nuclei. Still other thermonuclear reactions, such as helium fusion, carbon fusion, and oxygen fusion, occur late in the lives of many stars.

## 16-2 A theoretical model of the Sun shows how energy gets from its center to its surface

TUTORIAL 16.2 While thermonuclear fusion is the source of the Sun's energy, this process cannot take place everywhere within the Sun. As we have seen, extremely high temperatures—in excess of $10^7$ K—are required for atomic nuclei to fuse together to form larger nuclei. The temperature of the Sun's visible surface, about 5800 K, is too low for these reactions to occur there. Hence, thermonuclear fusion can be taking place only within the Sun's interior. But precisely where does it take place? And how does the energy produced by fusion make its way to the surface, where it is emitted into space in the form of photons?

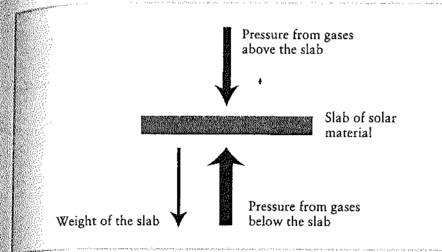
To answer these questions, we must understand conditions in the Sun's interior. Ideally, we would send an exploratory spacecraft to probe deep into the Sun; in practice, the Sun's intense heat would vaporize even the sturdiest spacecraft. Instead, astronomers use the laws of physics to construct a theoretical model of the Sun. (We discussed the use of models in science in Section 1-1.) Let's see what ingredients go into building a model of this kind.

### Hydrostatic Equilibrium

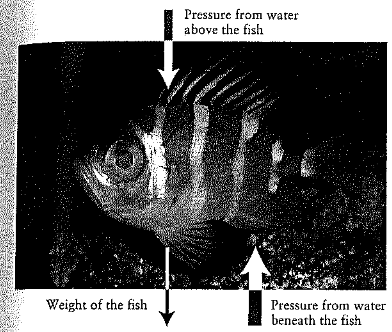
Note first that the Sun is not undergoing any dramatic changes. The Sun is not exploding or collapsing, nor is it significantly heating or cooling. The Sun is thus said to be in both *hydrostatic equilibrium* and *thermal equilibrium*.

To understand what is meant by **hydrostatic equilibrium,** imagine a slab of material in the solar interior (Figure 16-2*a*). In equilibrium, the slab on average will move neither up nor down. (In fact, there are upward and downward motions of material inside the Sun, but these motions average out.) Equilibrium is maintained by a balance among three forces that act on this slab:

1. The downward pressure of the layers of solar material above the slab.
2. The upward pressure of the hot gases beneath the slab.
3. The slab's weight—that is, the downward gravitational pull it feels from the rest of the Sun.



(a) Material inside the sun is in hydrostatic equilibrium, so forces balance



(b) A fish floating in water is in hydrostatic equilibrium, so forces balance

**Figure 16-2**

Hydrostatic Equilibrium **(a)** Material in the Sun's interior tends to move neither up nor down. The upward forces on a slab of solar material (due to pressure of gases below the slab) must balance the downward forces (due to the slab's weight and the pressure of gases above the slab). Hence, the pressure must increase with increasing depth. **(b)** The same principle applies to a fish floating in water. In equilibrium, the forces balance and the fish neither rises nor sinks. (Ken Usami/PhotoDisc)

The pressure from below must balance both the slab's weight and the pressure from above. Hence, the pressure below the slab must be greater than that above the slab. In other words, pressure has to increase with increasing depth. For the same reason, pressure increases as you dive deeper into the ocean (Figure 16-2*b*) or as you move toward lower altitudes in our atmosphere.

Hydrostatic equilibrium also tells us about the density of the slab. If the slab is too dense, its weight will be too large and it will sink; if the density is too low, the slab will rise. To prevent this, the density of solar material must have a certain value at each depth within the solar interior. (The same principle applies to objects that float beneath the surface of the ocean. Scuba divers wear weight belts to increase their average density so that they will neither rise nor sink but will stay submerged at the same level.)

### Thermal Equilibrium

Another consideration is that the Sun's interior is so hot that it is completely gaseous. Gases compress and become more dense when you apply greater pressure to them, so density must increase along with pressure as you go to greater depths within the Sun. Furthermore, when you compress a gas, its temperature tends to increase, so the temperature must also increase as you move toward the Sun's center.

While the temperature in the solar interior is different at different depths, the temperature at each depth remains constant in time. This principle is called **thermal equilibrium.** For the Sun to be in thermal equilibrium, all the energy generated by thermonuclear reactions in the Sun's core must be transported to the Sun's glowing surface, where it can be radiated into space. If too much energy flowed from the core to the surface to be radiated away, the Sun's interior would cool down; alternatively, the Sun's interior would heat up if too little energy flowed to the surface.

### Transporting Energy Outward from the Sun's Core

But exactly how is energy transported from the Sun's center to its surface? There are three methods of energy transport: *conduction, convection,* and *radiative diffusion.* Only the last two are important inside the Sun.

If you heat one end of a metal bar with a blowtorch, energy flows to the other end of the bar so that it too becomes warm. The efficiency of this method of energy transport, called **conduction,** varies significantly from one substance to another. For example, metal is a good conductor of heat, but plastic is not (which is why metal cooking pots often have plastic handles). Conduction is not an efficient means of energy transport in substances with low average densities, including the gases inside stars like the Sun.

Inside stars like our Sun, energy moves from center to surface by two other means: convection and radiative diffusion. **Convection** is the circulation of fluids—gases or liquids—between hot and cool regions. Hot gases (with lower density) rise toward a star's surface, while cool gases (with higher density) sink back down toward the star's center. This physical movement of gases transports heat energy outward in a star, just as the physical movement of water boiling in a pot transports energy from the bottom of the pot (where the heat is applied) to the cooler water at the surface.

In **radiative diffusion,** photons created in the thermonuclear inferno at a star's center diffuse outward toward the star's surface. Individual photons are absorbed and reemitted by atoms and electrons inside the star. The overall result is an outward migration from the hot core, where photons are constantly created, toward the cooler surface, where they escape into space.

### Modeling the Sun

To construct a model of a star like the Sun, astrophysicists express the ideas of hydrostatic equilibrium, thermal equilibrium, and energy transport as a set of equations. To ensure that the model applies to the particular star under study, they also make

## Table 16-2 A Theoretical Model of the Sun

| Distance from the Sun's center (solar radii) | Fraction of luminosity | Fraction of mass | Temperature ($\times 10^6$ K) | Density (kg/m$^3$) | Pressure relative to pressure at center |
|---|---|---|---|---|---|
| 0.0 | 0.00 | 0.00 | 15.5 | 160,000 | 1.00 |
| 0.1 | 0.42 | 0.07 | 13.0 | 90,000 | 0.46 |
| 0.2 | 0.94 | 0.35 | 9.5 | 40,000 | 0.15 |
| 0.3 | 1.00 | 0.64 | 6.7 | 13,000 | 0.04 |
| 0.4 | 1.00 | 0.85 | 4.8 | 4,000 | 0.007 |
| 0.5 | 1.00 | 0.94 | 3.4 | 1,000 | 0.001 |
| 0.6 | 1.00 | 0.98 | 2.2 | 400 | 0.0003 |
| 0.7 | 1.00 | 0.99 | 1.2 | 80 | $4 \times 10^{-5}$ |
| 0.8 | 1.00 | 1.00 | 0.7 | 20 | $5 \times 10^{-6}$ |
| 0.9 | 1.00 | 1.00 | 0.3 | 2 | $3 \times 10^{-7}$ |
| 1.0 | 1.00 | 1.00 | 0.006 | 0.00030 | $4 \times 10^{-13}$ |

*Note: The distance from the Sun's center is expressed as a fraction of the Sun's radius ($R_\odot$). Thus, 0.0 is at the center of the Sun and 1.0 is at the surface. The fraction of luminosity is that portion of the Sun's total luminosity produced within each distance from the center; this is equal to 1.00 for distances of 0.25 $R_\odot$ or more, which means that all of the Sun's nuclear reactions occur within 0.25 solar radius from the Sun's center. The fraction of mass is that portion of the Sun's total mass lying within each distance from the Sun's center. The pressure is expressed as a fraction of the pressure at the center of the Sun.*

use of astronomical observations of the star's surface. (For example, to construct a model of the Sun, they use the data that the Sun's surface temperature is 5800 K, its luminosity is $3.9 \times 10^{26}$ W, and the gas pressure and density at the surface are almost zero.) The astrophysicists then use a computer to solve their set of equations and calculate conditions layer by layer in toward the star's center. The result is a model of how temperature, pressure, and density increase with increasing depth below the star's surface.

Table 16-2 and Figure 16-3 show a theoretical model of the Sun that was calculated in just this way. Different models of the Sun use slightly different assumptions, but all models give essentially the same results as those shown here. From such computer models we have learned that at the Sun's center the density is 160,000 kg/m$^3$ (14 times the density of lead!), the temperature is $1.55 \times 10^7$ K, and the pressure is $3.4 \times 10^{11}$ atm. (One atmosphere, or 1 atm, is the average atmospheric pressure at sea level on Earth.)

Table 16-2 and Figure 16-3 show that the solar luminosity rises to 100% at about one-quarter of the way from the Sun's center to its surface. In other words, the Sun's energy production occurs within a volume that extends out only to 0.25 $R_\odot$. (The symbol $R_\odot$ denotes the solar radius, or radius of the Sun as a whole, equal to 696,000 km.) Outside 0.25 $R_\odot$, the density and temperature are too low for thermonuclear reactions to take place. Also note that 94% of the total mass of the Sun is found within the inner 0.5 $R_\odot$. Hence, the outer 0.5 $R_\odot$ contains only a relatively small amount of material.

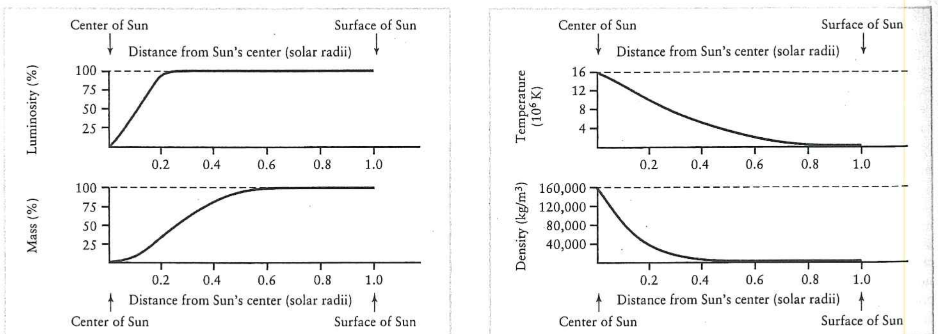
How energy flows from the Sun's center toward its surface depends on how easily photons move through the gas. If the solar gases are comparatively transparent, photons can travel moderate distances before being scattered or absorbed, and energy is thus transported by radiative diffusion. If the gases are comparatively *opaque*, photons are frequently scattered or absorbed and can't easily get through the gas. In an opaque gas, heat builds up and convection then becomes the most efficient means of energy transport. The gases start to churn, with hot lower-density moving upward and cooler gas sinking downward.

From the center of the Sun out to about 0.71 $R_\odot$, energy is transported by radiative diffusion. Hence, this region is called the **radiative zone.** Beyond about 0.71 $R_\odot$, the temperature is low enough (a mere $2 \times 10^6$ K or so) for electrons and hydrogen nuclei to join into hydrogen atoms. These atoms are very effective at absorbing photons, much more so than free electrons or nuclei, and this absorption chokes off the outward flow of photons. Therefore, beyond about 0.71 $R_\odot$, radiative diffusion is not an effective way to transport energy. Instead, convection dominates the energy flow in this outer region, which is why it is called the **convective zone.** Figure 16-4 shows these aspects of the Sun's internal structure.

Although energy travels through the radiative zone in the form of photons, the photons have a difficult time of it. Table 16-2 shows that the material in this zone is extremely dense, so photons from the Sun's core take a long time to diffuse through the radiative zone. As a result, it takes approximately 170,000 years for energy created at the Sun's center to travel 696,000 km to the solar surface and finally escape as sunlight. The energy flows outward at an average rate of 50 centimeters per hour, or about 20 times slower than a snail's pace.

Once the energy escapes from the Sun, it travels much faster—at the speed of light. Thus, solar energy that reaches you today took only 8 minutes to travel the 150 million kilometers from the Sun's surface to Earth. But this energy was actually produced by thermonuclear reactions that took place in the Sun's core hundreds of thousands of years ago.

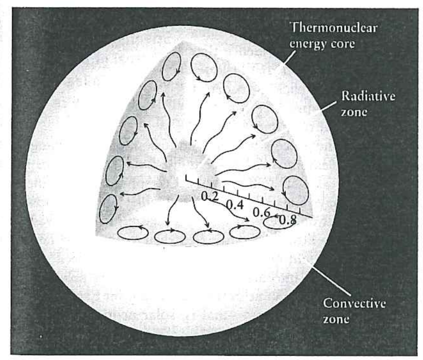
**The sunlight that reaches Earth today results from thermonuclear reactions that took place about 170,000 years ago**




## Figure 16-3

INTERACTIVE EXERCISE 16.1

A Theoretical Model of the Sun's Interior These graphs depict what percentage of the Sun's total luminosity is produced within each distance from the center (upper left), what percentage of the total mass lies within each distance from the center (lower left), the temperature at each distance (upper right), and the density at each distance (lower right). (See Table 16-2 for a numerical version of this model.)




## Figure 16-4

The Sun's Internal Structure Thermonuclear reactions occur in the Sun's core, which extends out to a distance of 0.25 $R_\odot$ from the center. Energy is transported outward, via radiative diffusion, to a distance of about 0.71 $R_\odot$. In the outer layers between 0.71 $R_\odot$ and 1.00 $R_\odot$, energy flows outward by convection.

## 16-3 Astronomers probe the solar interior using the Sun's own vibrations

WEB LINK 16.2

We have described how astrophysicists construct models of the Sun. But since we cannot see into the Sun's opaque interior, how can we check these models to see if they are accurate? What is needed is a technique for probing the Sun's interior. A very powerful technique of just this kind involves measuring vibrations of the Sun as a whole. This field of solar research is called **helioseismology.**

Vibrations are a useful tool for examining the hidden interiors of all kinds of objects. Food shoppers test whether melons are ripe by tapping on them and listening to the vibrations. Geologists can determine the structure of Earth's interior by using seismographs to record vibrations during earthquakes.

Although there are no true "sunquakes," the Sun does vibrate at a variety of frequencies, somewhat like a ringing bell. These vibrations were first noticed in 1960 by Robert Leighton of the California Institute of Technology, who made high-precision Doppler shift observations of the solar surface. These measurements revealed that parts of the Sun's surface move up and down about 10 meters every 5 minutes. Since the mid-1970s, several astronomers have reported slower vibrations, having periods ranging from 20 to 160 minutes. The detection of extremely slow vibrations has inspired astronomers to organize networks of telescopes around and in orbit above Earth to monitor the Sun's vibrations on a continuous basis.

WEB LINK 16.3

The vibrations of the Sun's surface can be compared with sound waves. If you could somehow survive within the Sun's outermost layers, you would first notice a deafening roar, somewhat like a jet engine, produced by

## Activities

### Observing Projects

Observing tips and tools

At the risk of repeating ourselves, we remind you to ***never look directly at the Sun, because it can easily cause permanent blindness.*** You can view the Sun safely without a telescope just by using two pieces of white cardboard. First, use a pin to poke a small hole in one piece of cardboard; this will be your "lens," and the other piece of cardboard will be your "viewing screen." Hold the "lens" piece of cardboard so that it is face-on to the Sun and sunlight can pass through the hole. With your other hand, hold the "viewing screen" so that the sunlight from the "lens" falls on it. Adjust the distance between the two pieces of cardboard so that you see a sharp image of the Sun on the "viewing screen." This image is perfectly safe to view. It is actually possible to see sunspots with this low-tech apparatus.

For a better view, use a telescope with a solar filter that fits on the front of the telescope. A standard solar filter is a piece of glass coated with a thin layer of metal to give it a mirrorlike appearance. This coating reflects almost all the sunlight that falls on it, so that only a tiny, safe amount of sunlight enters the telescope. An $H_\alpha$ filter, which looks like a red piece of glass, keeps the light at a safe level by admitting only a very narrow range of wavelengths. (Filters that fit on the back of the telescope are *not* recommended. The telescope focuses concentrated sunlight on such a filter, heating it and making it susceptible to cracking—and if the filter cracks when you are looking through it, your eye will be ruined instantly and permanently.)

To use a telescope with a solar filter, first aim the telescope away from the Sun, then put on the filter. Keep the lens cap on the telescope's secondary wide-angle "finder scope" (if it has one), because the heat of sunlight can fry the finder scope's optics. Next, aim the telescope toward the Sun, using the telescope's shadow to judge when you are pointed in the right direction. You can then safely look through the telescope's eyepiece. When you are done, make sure you point the telescope away from the Sun before removing the filter and storing the telescope.

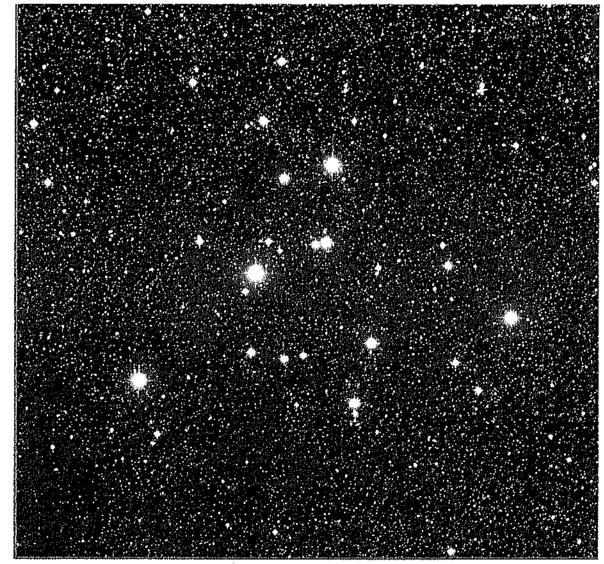
Note that the amount of solar activity that you can see (sunspots, filaments, flares, prominences, and so on) will depend on where the Sun is in its 11-year sunspot cycle.

58. Use a telescope with a solar filter to observe the surface of the Sun. Do you see any sunspots? Sketch their appearance. Can you distinguish between the umbrae and penumbrae of the sunspots? Can you see limb darkening? Can you see any granulation?
59. If you have access to an $H_\alpha$ filter attached to a telescope especially designed for viewing the Sun safely, use this instrument to examine the solar surface. How does the appearance of the Sun differ from that in white light? What do sunspots look like in $H_\alpha$? Can you see any prominences? Can you see any filaments? Are the filaments in the $H_\alpha$ image near any sunspots seen in white light? (Note that the amount of activity that you see will be much greater at some times during the solar cycle than at others.)
60. Use the *Starry Night Enthusiast*™ program to measure the Sun's rotation. Display the entire celestial sphere by selecting **Guides > Atlas** in the **Favorites** menu and center on the Sun by double-clicking on **Sun** in the **Find** pane. Using the controls at the right-hand end of the toolbar, zoom in until you can see details on the Sun's surface clearly. In the toolbar, set the **Time Flow Rate** to 1 day. Using the time forward and backward buttons in the toolbar, step through enough time to determine the rotation period of the Sun. Which part of the actual Sun's surface rotates at the rate shown in *Starry Night Enthusiast*™? (Note: The program does not show the Sun's differential rotation.)
61. Use the *Starry Night Enthusiast*™ program to examine the Sun. Open the **Favorites** pane and double-click on **Solar System > Inner Solar System** to display the inner planets surrounding the Sun. Stop **Time Advance** and use the down arrow in the toolbar under **Viewing Location** to approach to within about 0.015 AU of the Sun. You can rotate the Sun by placing the mouse cursor over the image and, while holding down the **Shift** key, hold down the mouse button while moving the mouse. (On a two-button mouse, hold down the left mouse button.) (a) The Sun's equator lies close to the plane of the ecliptic. Where do most of the sunspots visible on the image of the Sun lie relative to the solar equator? Check Figure 16-18 for more realistic images of sunspots and Figure 16-19 for the latitude distribution of sunspots. (b) Based on your observations in (a), does the image in *Starry Night Enthusiast*™ show the Sun near the beginning, middle, or end of the 11-year sunspot cycle? Explain your reasoning. You can see current solar images from both ground and space-based solar telescopes by opening the LiveSky pane if you have an Internet connection on your computer.

### Collaborative Exercises

62. Figure 16-19 shows variations in the average latitude of sunspots. Estimate the average latitude of sunspots in the year you were born and estimate the average latitude on your twenty-first birthday. Make rough sketches of the Sun during those years to illustrate your answers.
63. Create a diagram showing a sketch of how limb darkening on the Sun would look different if the Sun had either a thicker or thinner photosphere. Be sure to include a caption explaining your diagram.
64. Solar granules, shown in Figure 16-9, are about 1000 km across. What city is about that distance away from where you are right now? What city is that distance from the birthplace of each group member?
65. Magnetic arches in the corona are shown in Figure 16-25*a*. How many Earths high are these arches, and how many Earths could fit inside one arch?

# 17 The Nature of the Stars



R I V U X G

Some stars in this cluster (called M39) are distinctly blue in color, while others are yellow or red. (Heidi Schweiker/NOAO/AURA/NSF)

To the unaided eye, the night sky is spangled with several thousand stars, each appearing as a bright pinpoint of light. With a pair of binoculars, you can see some 10,000 other, fainter stars; with a 15-cm (6-in) telescope, the total rises to more than 2 million. Astronomers now know that there are in excess of 100 billion ($10^{11}$) stars in our Milky Way Galaxy alone.

But what are these distant pinpoints? To the great thinkers of ancient Greece, the stars were bits of light embedded in a vast sphere with Earth at the center. They thought the stars were composed of a mysterious "fifth element," quite unlike anything found on Earth.

Today, we know that the stars are made of the same chemical elements found on Earth. We know their sizes, their temperatures, their masses, and something of their internal structures. We understand, too, why the stars in the accompanying image come in a range of beautiful colors: Blue stars have high surface temperatures, while the surface temperatures of red and yellow stars are relatively low.

How have we learned these things? How can we know the nature of the stars, objects so distant that their light takes years or centuries to reach us? In this chapter, we will learn about the measurements and calculations that astronomers make to determine the properties of stars. We will also take a first look at the Hertzsprung-Russell diagram, an important tool that helps astronomers systematize the wealth of available information about the stars. In later chapters, we will use this diagram to help us understand how stars are born, evolve, and eventually die.

## Learning Goals

*By reading the sections of this chapter, you will learn*

17-1 How we can measure the distances to the stars
17-2 How we measure a star's brightness and luminosity
17-3 The magnitude scale for brightness and luminosity
17-4 How a star's color indicates its temperature
17-5 How a star's spectrum reveals its chemical composition
17-6 How we can determine the sizes of stars
17-7 How H-R diagrams summarize our knowledge of the stars
17-8 How we can deduce a star's size from its spectrum
17-9 How we can use binary stars to measure the masses of stars
17-10 How we can learn about binary stars in very close orbits
17-11 What eclipsing binaries are and what they tell us about the sizes of stars

## 17-1 Careful measurements of the parallaxes of stars reveal their distances

The vast majority of stars are objects very much like the Sun. This understanding followed from the discovery that the stars are tremendously far from us, at distances so great that their light takes years to reach us. Because the stars at night are clearly visible to the naked eye despite these huge distances, it must be that the **luminosity** of the stars—that is, how much energy they emit into space per second—is comparable to or greater than that of the Sun. Just as for the Sun, the only explanation for such tremendous luminosities is that thermonuclear reactions are occurring within the stars (see Section 16-1).

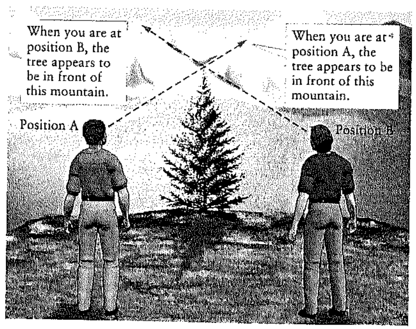
Clearly, then, it's important to know how distant the stars are. But how do we measure these distances? You might think these distances are determined by comparing how bright different stars appear. Perhaps the star Betelgeuse in the constellation Orion appears bright because it is relatively close, while the dimmer and less conspicuous star Polaris (the North Star, in the constellation Ursa Minor) is farther away.

But this line of reasoning is incorrect: Polaris is actually closer to us than Betelgeuse! How bright a star appears is *not* a good indicator of its distance. If you see a light on a darkened road, it could be a motorcycle headlight a kilometer away or a person holding a flashlight just a few meters away. In the same way, a bright star might be extremely far away but have an unusually high luminosity, and a dim star might be relatively close but have a rather low luminosity. Astronomers must use other techniques to determine the distances to the stars.

## Parallax and the Distances to the Stars

TUTORIAL 17.1

The most straightforward way of measuring stellar distances uses an effect called **parallax**, which is the apparent displacement of an object because of a change in the observer's point of view (Figure 17-1). To see how parallax works, hold your arm out straight in front of you. Now look at the hand on your outstretched arm, first with your left eye closed, then with your right eye closed. When you close one eye and open the other, your hand appears to shift back and forth against the background of more distant objects.




### Figure 17-1

Parallax Imagine looking at some nearby object (a tree) against a distant background (mountains). When you move from one location to another, the nearby object appears to shift with respect to the distant background scenery. This familiar phenomenon is called parallax.

The closer the object you are viewing, the greater the parallax shift. To see this increased shift, repeat the experiment with your hand held closer to your face. Your brain analyzes such parallax shifts constantly as it compares the images from your left and right eyes, and in this way determines the distances to objects around you. This analysis is the basis for depth perception.

**You measure distances around you by comparing the images from your left and right eyes—we find the distances to stars using the same principle**

To measure the distance to a star, astronomers measure the parallax shift of the star using two points of view that are as far apart as possible—at opposite sides of Earth's orbit. The direction from Earth to a nearby star changes as our planet orbits the Sun, and the nearby star appears to move back and forth against the background of more distant stars (Figure 17-2). This motion is called **stellar parallax**. The parallax ($p$) of a star is equal to half the angle through which the star's apparent position shifts as Earth moves from one side of its orbit to the other. The larger the parallax $p$, the smaller the distance $d$ to the star (compare Figure 17-2*a* with Figure 17-2*b*).

It is convenient to measure the distance $d$ in **parsecs**. A star with a parallax angle of 1 second of arc ($p$ = 1 arcsec) is at a distance of 1 parsec ($d$ = 1 pc). (The word "parsec" is a contraction of the phrase "the distance at which a star has a *par*allax of one arc*sec*ond." Recall from Section 1-7 that 1 parsec equals 3.26 light-years, $3.09 \times 10^{13}$ km, or 206,265 AU; see Figure 1-14.) If the angle $p$ is measured in arcseconds, then the distance $d$ to the star in parsecs is given by the following equation:

**Relation between a star's distance and its parallax**

$$d = \frac{1}{p}$$

$d$ = distance to a star, in parsecs

$p$ = parallax angle of that star, in arcseconds

This simple relationship between parallax and distance in parsecs is one of the main reasons that astronomers usually measure cosmic distances in parsecs rather than light-years. For example, a star whose parallax is $p$ = 0.1 arcsec is at a distance $d$ = 1/(0.1) = 10 parsecs from Earth. Barnard's star, named for the American astronomer Edward E. Barnard, has a parallax of 0.547 arcsec. Hence, the distance to this star is

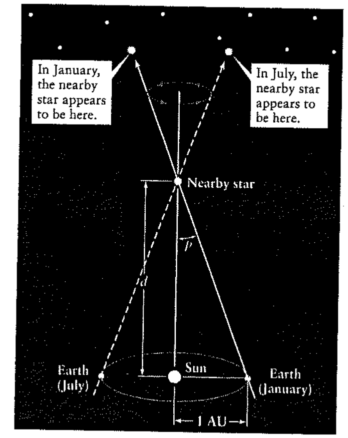
$$d = \frac{1}{p} = \frac{1}{0.547} = 1.83 \text{ pc}$$

Because 1 parsec is 3.26 light-years, this distance can also be expressed as

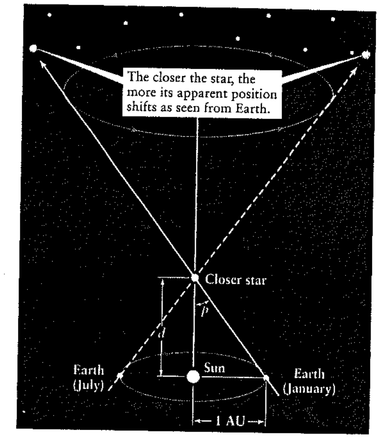
$$d = 1.83 \text{ pc} \times \frac{3.26 \text{ ly}}{1 \text{ pc}} = 5.97 \text{ ly}$$

All known stars have parallax angles less than one arcsecond. In other words, the closest star is more than 1 parsec away. Such






(a) Parallax of a nearby star




(b) Parallax of an even closer star

ANIMATION 17.1

### Figure 17-2

Stellar Parallax **(a)** As Earth orbits the Sun, a nearby star appears to shift its position against the background of distant stars. The parallax ($p$) of the star is equal to the angular radius of Earth's orbit as seen from the star. **(b)** The closer the star is to us, the greater the parallax angle $p$. The distance $d$ to the star (in parsecs) is equal to the reciprocal of the parallax angle $p$ (in arcseconds): $d = 1/p$.

small parallax angles are difficult to detect, so it was not until 1838 that the first successful parallax measurements were made by the German astronomer and mathematician Friedrich Wilhelm Bessel. He found the parallax angle of the star 61 Cygni to be just 1/3 arcsec—equal to the angular diameter of a dime at a distance of 11 kilometers, or 7 miles. He thus determined that this star is about 3 pc from Earth. (Modern measurements give a slightly smaller parallax angle, which means that 61 Cygni is actually more than 3 pc away.) The star Proxima Centauri has the largest known parallax angle, 0.772 arcsec, and hence is the closest known star (other than the Sun); its distance is 1/(0.772) = 1.30 pc.

WEB LINK 17.3 AIMM 17.1

Appendix 4 at the back of this book lists all the stars within 4 pc of the Sun, as determined by parallax measurements. Most of these stars are far too dim to be seen with the naked eye, which is why their names are probably unfamiliar to you. By contrast, the majority of the familiar, bright stars in the nighttime sky (listed in Appendix 5) are so far away that their parallaxes cannot be measured from Earth's surface. They appear bright not because they are close, but because they are far more luminous than the Sun. The brightest stars in the sky are *not* necessarily the nearest stars!

## Measuring Parallax from Space

Parallax angles smaller than about 0.01 arcsec are extremely difficult to measure from Earth, in part because of the blurring effects of the atmosphere. Therefore, the parallax method used with ground-based telescopes can give fairly reliable distances only for stars nearer than about 1/0.01 = 100 pc. But an observatory in space is unhampered by the atmosphere. Observations made from spacecraft therefore permit astronomers to measure even smaller parallax angles and thus determine the distances to more remote stars.

WEB LINK 17.4

In 1989 the European Space Agency (ESA) launched the satellite *Hipparcos,* an acronym for *H*igh *P*recision *Par*allax *Co*llecting *S*atellite (and a commemoration of the ancient Greek astronomer Hipparchus, who created one of the first star charts). Over more than three years of observations, the telescope aboard *Hipparcos* was used to measure the parallaxes of 118,000 stars, some with an accuracy of 0.001 arcsecond. This telescope has enabled astronomers to determine stellar distances out to several hundred parsecs, and with much greater precision than has been possible with ground-based observations. In the years to come, astronomers will increasingly turn to space-based observations to determine stellar distances.

Unfortunately, most of the stars in the Galaxy are so far away that their parallax angles are too small to measure even with an orbiting telescope. Later in this chapter, we will discuss a technique that can be used to find the distances to these more remote stars. In Chapters 24 and 26 we will learn about other techniques that astronomers use to determine the much larger distances to galaxies beyond the Milky Way. These techniques also

## BOX 17-1 Tools of the Astronomer's Trade

### Stellar Motions

Stars can move through space in any direction. The **space velocity** of a star describes how fast and in what direction it is moving. As the accompanying figure shows, a star's space velocity $v$ can be broken into components parallel and perpendicular to our line of sight.

The component perpendicular to our line of sight—that is, across the plane of the sky—is called the star's **tangential velocity** ($v_t$). To determine it, astronomers must know the distance to a star ($d$) and its **proper motion** ($\mu$, the Greek letter mu), which is the number of arcseconds that the star appears to move per year on the celestial sphere. Proper motion does not repeat itself yearly, so it can be distinguished from the apparent back-and-forth motion due to parallax. In terms of a star's distance and proper motion, its tangential velocity (in km/s) is

$$v_t = 4.74\mu d$$

where $\mu$ is in arcseconds per year and $d$ is in parsecs. For example, Barnard's star (Figure 17-3) has a proper motion of 10.358 arcseconds per year and a distance of 1.83 pc. Hence, its tangential velocity is

$$v_t = 4.74(10.358)(1.83) = 89.8 \text{ km/s}$$

The component of a star's motion parallel to our line of sight—that is, either directly toward us or directly away from us—is its **radial velocity** ($v_r$). It can be determined from measurements of the Doppler shifts of the star's spectral lines (see Section 5-9 and Box 5-6). If a star is approaching us, the wavelengths of all of its spectral lines are decreased (blueshifted); if the star is receding from us, the wavelengths are increased (redshifted). The radial velocity $v_r$ is related to the wavelength shift by the equation

$$\frac{\lambda - \lambda_0}{\lambda_0} = \frac{v_r}{c}$$

In this equation, $\lambda$ is the wavelength of light coming from the star, $\lambda_0$ is what the wavelength would be if the star were not moving, and $c$ is the speed of light. As an illustration, a particular spectral line of iron in the spectrum of Barnard's star has a wavelength ($\lambda$) of 516.445 nm. As measured in a laboratory on Earth, the same spectral line has a wavelength ($\lambda_0$) of 516.629 nm. Thus, for Barnard's star, our equation becomes

$$\frac{516.445 \text{ nm} - 516.629 \text{ nm}}{516.629 \text{ nm}} = -0.000356 = \frac{v_r}{c}$$

Solving this equation for the radial velocity $v_r$, we find

$$v_r = (-0.000356)\, c = (-0.000356)(3.00 \times 10^5 \text{ km/s}) = -107 \text{ km/s}$$

The minus sign means that Barnard's star is moving toward us. You can check this interpretation by noting that the wavelength $\lambda = 516.445$ nm received from Barnard's star is less than the laboratory wavelength $\lambda_0 = 516.629$ nm; hence, the light from the star is blueshifted, which indeed means that the star is approaching. If the star were receding, its light would be redshifted, and its radial velocity would be positive.

The illustration shows that the tangential velocity and radial velocity form two sides of a right triangle. The long side (hypotenuse) of this triangle is the space velocity ($v$). From the Pythagorean theorem, the space velocity is

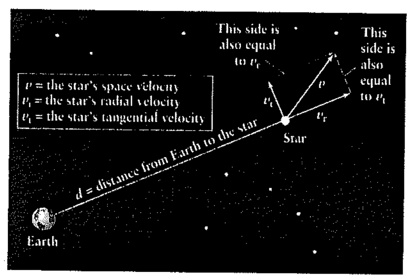
$$v = \sqrt{v_t^2 + v_r^2}$$

For Barnard's star, the space velocity is

$$v = \sqrt{(89.4 \text{ km/s})^2 + (-107 \text{ km/s})^2} = 140 \text{ km/s}$$

Therefore, Barnard's star is moving through space at a speed of 140 km/s (503,000 km/h, or 312,000 mi/h) relative to the Sun.

Determining the space velocities of stars is essential for understanding the structure of the Galaxy. Studies show that the stars in our local neighborhood are moving in wide orbits around the center of the Galaxy, which lies some 8000 pc (26,000 light-years) away in the direction of the constellation Sagittarius (the Archer). While many of the orbits are roughly circular and lie in nearly the same plane, others are highly elliptical or steeply inclined to the galactic plane. We will see in Chapter 23 how the orbits of stars and gas clouds reveal the Galaxy's spiral structure.



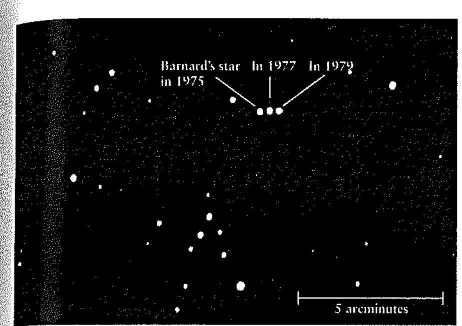




**Figure 17-3** R I V U X G
The Motion of Barnard's Star Three photographs taken over a four-year period were combined to show the motion of Barnard's star, which lies 1.82 parsecs away in the constellation Ophiuchus. Over this time interval, Barnard's star moved more than 41 arcseconds on the celestial sphere (about 0.69 arcminutes, or 0.012°), more than any other star. (John Sanford/Science Photo Library)

help us understand the overall size, age, and structure of the universe.

### The Importance of Parallax Measurements

Because it can be used only on relatively close stars, stellar parallax might seem to be of limited usefulness. But parallax measurements are the cornerstone for all other methods of finding the distances to remote objects. These other methods require a precise and accurate knowledge of the distances to nearby stars, as determined by stellar parallax. Hence, any inaccuracies in the parallax angles for nearby stars can translate into substantial errors in measurement for the whole universe. To minimize these errors astronomers are continually trying to perfect their parallax-measuring techniques.

Stellar parallax is an *apparent* motion of stars caused by Earth's orbital motion around the Sun. But stars are not fixed objects and actually do move through space. As a result, stars change their positions on the celestial sphere (Figure 17-3), and they move either toward or away from the Sun. These motions are sufficiently slow, however, that changes in the positions of the stars are hardly noticeable over a human lifetime. Box 17-1 describes how astronomers study these motions and what insights they gain from these studies.

## 17-2 If a star's distance is known, its luminosity can be determined from its apparent brightness

All the stars you can see in the nighttime sky shine by thermonuclear fusion, just as the Sun does (see Section 16-1). But they are by no means merely identical copies of the Sun. Stars differ in their luminosity ($L$), the amount of light energy they emit each second. Luminosity is usually measured either in watts (1 watt, or 1 W, is 1 joule per second) or as a multiple of the Sun's luminosity ($L_\odot$, equal to $3.90 \times 10^{26}$ W). Most stars are less luminous than the Sun, but some blaze forth with a million times the Sun's luminosity. Knowing a star's luminosity is essential for determining the star's history, its present-day internal structure, and its future evolution.

### Luminosity, Apparent Brightness, and the Inverse-Square Law

**Apparent brightness is a measure of how faint a star looks to us; luminosity is a measure of the star's total light output**

To determine the luminosity of a star, we first note that as light energy moves away from its source, it spreads out over increasingly larger regions of space. Imagine a sphere of radius $d$ centered on the light source, as in Figure 17-4. The amount of energy that passes each second through a square meter of the sphere's surface area is the total luminosity of the source ($L$) divided by the total surface area of the sphere (equal to $4\pi d^2$). This quantity is called the **apparent brightness** of the light, or just **brightness** ($b$), because how bright a light source appears depends on how much light energy per second enters through the area of a light detector (such as your eye). Apparent brightness is measured in watts per square meter (W/m$^2$). Written in the form of an equation, the relationship between apparent brightness and luminosity is

**Inverse-square law relating apparent brightness and luminosity**

$$b = \frac{L}{4\pi d^2}$$

$b$ = apparent brightness of a star's light, in W/m$^2$

$L$ = star's luminosity, in watts

$d$ = distance to star, in meters

This relationship is called the **inverse-square law,** because the apparent brightness of light that an observer can see or measure is inversely proportional to the square of the observer's distance ($d$) from the source. If you double your distance from a light source, its radiation is spread out over an area 4 times larger, so the apparent brightness you see is decreased by a factor of 4. Similarly, at triple the distance, the apparent brightness is 1/9 as great (see Figure 17-4).

We can apply the inverse-square law to the Sun, which is $1.50 \times 10^{11}$ m from Earth. Its apparent brightness ($b_\odot$) is

$$b_\odot = \frac{3.90 \times 10^{26} \text{ W}}{4\pi(1.50 \times 10^{11} \text{ m})^2} = 1370 \text{ W/m}^2$$

That is, a solar panel with an area of 1 square meter receives 1370 watts of power from the Sun.

With greater distance from the star, its light is spread over a larger area and its apparent brightness is less.

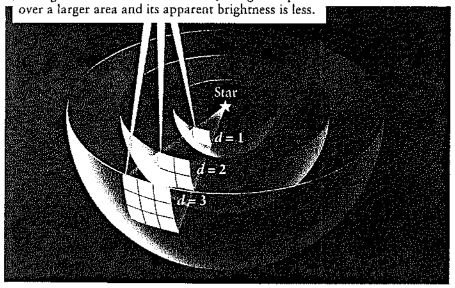



**Figure 17-4**

The Inverse-Square Law Radiation from a light source illuminates an area that increases as the square of the distance from the source. Hence, the apparent brightness decreases as the square of the distance. The brightness at $d = 2$ is $1/(2^2) = 1/4$ of the brightness at $d = 1$, and the brightness at $d = 3$ is $1/(3^2) = 1/9$ of that at $d = 1$.

Astronomers measure the apparent brightness of a star using a telescope with an attached light-sensitive instrument, similar to the light meter in a camera that determines the proper exposure. Measuring a star's apparent brightness is called **photometry.**

## Calculating a Star's Luminosity

The inverse-square law says that we can find a star's luminosity if we know its distance and its apparent brightness. For convenience, this law can be expressed in a somewhat different form. We first rearrange the above equation:

$$L = 4\pi d^2 b$$

We then apply this equation to the Sun. That is, we write a similar equation relating the Sun's luminosity ($L_\odot$), the distance from Earth to the Sun ($d_\odot$, equal to 1 AU), and the Sun's apparent brightness ($b_\odot$):

$$L_\odot = 4\pi d_\odot^2 b_\odot$$

If we take the ratio of these two equations, the unpleasant factor of $4\pi$ drops out and we are left with the following:

**Determining a star's luminosity from its apparent brightness**

$$\frac{L}{L_\odot} = \left(\frac{d}{d_\odot}\right)^2 \frac{b}{b_\odot}$$

$L/L_\odot$ = ratio of the star's luminosity to the Sun's luminosity

$d/d_\odot$ = ratio of the star's distance to the Earth-Sun distance

$b/b_\odot$ = ratio of the star's apparent brightness to the Sun's apparent brightness

We need to know just two things to find a star's luminosity: the distance to a star as compared to the Earth-Sun distance (the ratio $d/d_\odot$), and how that star's apparent brightness compares to that of the Sun (the ratio $b/b_\odot$). Then we can use the above equation to find how luminous that star is compared to the Sun (the ratio $L/L_\odot$).

In other words, this equation gives us a general rule relating the luminosity, distance, and apparent brightness of a star:

***We can determine the luminosity of a star from its distance and apparent brightness. For a given distance, the brighter the star, the more luminous that star must be. For a given apparent brightness, the more distant the star, the more luminous it must be to be seen at that distance.***

Box 17-2 shows how to use the above equation to determine the luminosity of the nearby star ε (epsilon) Eridani, the fifth brightest star in the constellation Eridanus (named for a river in Greek mythology). Parallax measurements indicate that ε Eridani is 3.23 pc away, and photometry shows that the star appears only $6.73 \times 10^{-13}$ as bright as the Sun. Using the above equation, we find that ε Eridani has only 0.30 times the luminosity of the Sun.

## The Stellar Population

Calculations of this kind show that stars come in a wide variety of different luminosities, with values that range from about $10^6 L_\odot$ (a million times the Sun's luminosity) to only about $10^{-4} L_\odot$ (a mere ten-thousandth of the Sun's light output). The most luminous star emits roughly $10^{10}$ times more energy each second than the least luminous! (To put this number in perspective, about $10^{10}$ human beings have lived on Earth since our species first evolved.)

As stars go, our Sun is neither extremely luminous nor extremely dim; it is a rather ordinary, garden-variety star. It is somewhat more luminous than most stars, however. Of more than 30 stars within 4 pc of the Sun (see Appendix 4), only three (α Centauri, Sirius, and Procyon) have a greater luminosity than the Sun.

To better characterize a typical population of stars, astronomers count the stars out to a certain distance from the Sun and plot the number of stars that have different luminosities. The resulting graph is called the **luminosity function.** Figure 17-5 on page 440 shows the luminosity function for stars in our part of the Milky Way Galaxy. The curve declines very steeply for the most luminous stars toward the left side of the graph, indicating that they are quite rare. For example, this graph shows that stars like the Sun are about 10,000 times more common than stars like Spica (which has a luminosity of 2100 $L_\odot$).

The exact shape of the curve in Figure 17-5 applies only to the vicinity of the Sun and similar regions in our Milky Way Galaxy. Other locations have somewhat different luminosity functions. In stellar populations in general, however, low-luminosity stars are much more common than high-luminosity ones.



BOX 17-2 Tools of the Astronomer's Trade

## Luminosity, Distance, and Apparent Brightness

The inverse-square law (Section 17-2) relates a star's luminosity, distance, and apparent brightness to the corresponding quantities for the Sun:

$$\frac{L}{L_\odot} = \left(\frac{d}{d_\odot}\right)^2 \frac{b}{b_\odot}$$

We can use a similar equation to relate the luminosities, distances, and apparent brightnesses of *any* two stars, which we call star 1 and star 2:

$$\frac{L_1}{L_2} = \left(\frac{d_1}{d_2}\right)^2 \frac{b_1}{b_2}$$

**EXAMPLE:** The star ε (epsilon) Eridani is 3.23 pc from Earth. As seen from Earth, this star appears only $6.73 \times 10^{-13}$ as bright as the Sun. What is the luminosity of ε Eridani compared with that of the Sun?

**Situation:** We are given the distance to ε Eridani ($d$ = 3.23 pc) and this star's brightness compared to that of the Sun ($b/b_\odot = 6.73 \times 10^{-13}$). Our goal is to find the ratio of the luminosity of ε Eridani to that of the Sun, that is, the quantity $L/L_\odot$.

**Tools:** Since we are asked to compare this star to the Sun, we use the first of the two equations given above, $L/L_\odot = (d/d_\odot)^2(b/b_\odot)$, to solve for $L/L_\odot$.

**Answer:** Our equation requires the ratio of the star's distance to the Sun's distance, $d/d_\odot$. The distance from Earth to the Sun is $d_\odot$ = 1 AU. To calculate the ratio $d/d_\odot$, we must express both distances in the same units. There are 206,265 AU in 1 parsec, so we can write the distance to ε Eridani as $d = (3.23\ \text{pc})(206{,}265\ \text{AU/pc}) = 6.66 \times 10^5$ AU. Hence, the ratio of distances is $d/d_\odot = (6.66 \times 10^5\ \text{AU})/(1\ \text{AU}) = 6.66 \times 10^5$. Then we find that the ratio of the luminosity of ε Eridani ($L$) to the Sun's luminosity ($L_\odot$) is

$$\frac{L}{L_\odot} = \left(\frac{d}{d_\odot}\right)^2 \frac{b}{b_\odot} = (6.66 \times 10^5)^2 \times (6.73 \times 10^{-13}) = 0.30$$

**Review:** This result means that ε Eridani is only 0.30 as luminous as the Sun; that is, its power output is only 30% as great.

**EXAMPLE:** Suppose star 1 is at half the distance of star 2 (that is, $d_1/d_2 = ½$) and that star 1 appears twice as bright as star 2 (that is, $b_1/b_2 = 2$). How do the luminosities of these two stars compare?

**Situation:** For these two stars, we are given the ratio of distances ($d_1/d_2$) and the ratio of apparent brightnesses ($b_1/b_2$). Our goal is to find the ratio of their luminosities ($L_1/L_2$).

**Tools:** Since we are comparing two stars, neither of which is the Sun, we use the second of the two equations above: $L_1/L_2 = (d_1/d_2)^2(b_1/b_2)$.

**Answer:** Plugging values into our equation, we find

$$\frac{L_1}{L_2} = \left(\frac{d_1}{d_2}\right)^2 \frac{b_1}{b_2} = \left(\frac{1}{2}\right)^2 \times 2 = \frac{1}{2}$$

**Review:** This result says that star 1 has only one-half the luminosity of star 2. Despite this, star 1 appears brighter than star 2 because it is closer to us.

The two equations above are also useful in the method of *spectroscopic parallax,* which we discuss in Section 17-8. It turns out that a star's luminosity can be determined simply by analyzing the star's spectrum. If the star's apparent brightness is also known, the star's distance can be calculated. The inverse-square law can be rewritten as an expression for the ratio of the star's distance from Earth ($d$) to the Earth-Sun distance ($d_\odot$):

$$\frac{d}{d_\odot} = \sqrt{\frac{(L/L_\odot)}{(b/b_\odot)}}$$

We can also use this formula as a relation between the properties of any two stars, 1 and 2:

$$\frac{d_1}{d_2} = \sqrt{\frac{(L_1/L_2)}{(b_1/b_2)}}$$

**EXAMPLE:** The star Pleione in the constellation Taurus is 190 times as luminous as the Sun but appears only $3.19 \times 10^{-13}$ as bright as the Sun. How far is Pleione from Earth?

**Situation:** We are told the ratio of Pleione's luminosity to that of the Sun ($L/L_\odot = 190$) and the ratio of their apparent brightnesses ($b/b_\odot = 3.19 \times 10^{-13}$). Our goal is to find the distance $d$ from Earth to Pleione.

**Tools:** Since we are comparing Pleione to the Sun, we use the first of the two equations above.

**Answer:** Our equation tells us the ratio of the Earth-Pleione distance to the Earth-Sun distance:

$$\frac{d}{d_\odot} = \sqrt{\frac{(L/L_\odot)}{(b/b_\odot)}} = \sqrt{\frac{190}{3.19 \times 10^{-13}}}$$

$$= \sqrt{5.95 \times 10^{14}} = 2.44 \times 10^7$$

*(continued on the next page)*

## BOX 17-2 (continued)

Hence, the distance from Earth to Pleione is $2.44 \times 10^7$ times greater than the distance from Earth to the Sun. The Sun-Earth distance is $d_\odot = 1$ AU and 206,265 AU = 1 pc, so we can express the star's distance as $d = (2.44 \times 10^7 \text{ AU}) \times (1 \text{ pc}/206{,}265 \text{ AU}) = 118$ pc.

**Review:** We can check our result by comparing it with the above example about the star ε Eridani. Pleione has a much greater luminosity than ε Eridani (190 times the Sun's luminosity versus 0.30 times), but Pleione appears dimmer than ε Eridani ($3.19 \times 10^{-13}$ times as bright as the Sun compared to $6.73 \times 10^{-13}$ times). For this to be true, Pleione must be much farther away from Earth than is ε Eridani. This is just what our results show: $d = 118$ pc for Pleione compared to $d = 3.23$ pc for ε Eridani.

**EXAMPLE:** The star δ (delta) Cephei, which lies 300 pc from Earth, is thousands of times more luminous than the Sun. Thanks to this great luminosity, stars like δ Cephei can be seen in galaxies millions of parsecs away. As an example, the Hubble Space Telescope has detected stars like δ Cephei within the galaxy NGC 3351, which lies in the direction of the constellation Leo. These stars appear only $9 \times 10^{-10}$ as bright as δ Cephei. What is the distance to NGC 3351?

**Situation:** To determine the distance we want, we need to find the distance to a star within NGC 3351. We are told that certain stars within this galaxy are identical to δ Cephei but appear only $9 \times 10^{-10}$ as bright.

**Tools:** We use the equation $d_1/d_2 = \sqrt{(L_1/L_2)/(b_1/b_2)}$ to relate two stars, one within NGC 3351 (call this star 1) and the identical star δ Cephei (star 2). Our goal is to find $d_1$.

**Answer:** Since the two stars are identical, they have the same luminosity ($L_1 = L_2$, or $L_1/L_2 = 1$). The brightness ratio is $b_1/b_2 = 9 \times 10^{-10}$, so our equation tells us that

$$\frac{d_1}{d_2} = \sqrt{\frac{(L_1/L_2)}{(b_1/b_2)}} = \sqrt{\frac{1}{9 \times 10^{-10}}} = \sqrt{1.1 \times 10^9} = 33{,}000$$

Hence, NGC 3351 is 33,000 times farther away than δ Cephei, which is 300 pc from Earth. The distance from Earth to NGC 3351 is therefore $(33{,}000)(300 \text{ pc}) = 10^7$ pc, or 10 megaparsecs (10 Mpc).

**Review:** This example illustrates one technique that astronomers use to measure extremely large distances. We will learn more about stars like δ Cephei in Chapter 19, and in Chapter 24 we will explore further how they are used to determine the distances to remote galaxies.

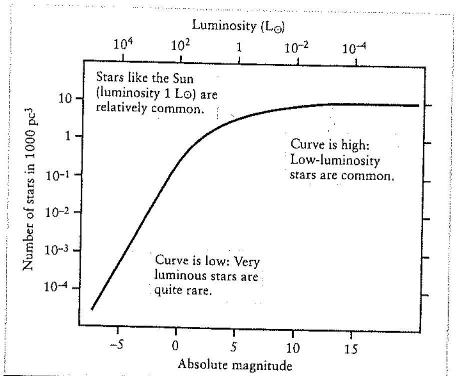



**Figure 17-5**

The Luminosity Function This graph shows how many stars of a given luminosity lie within a representative 1000 cubic-parsec volume. The scale at the bottom of the graph shows absolute magnitude, an alternative measure of a star's luminosity (described in Section 17-3). (Adapted from J. Bahcall and R. Soneira)

## 17-3 Astronomers often use the magnitude scale to denote brightness

Because astronomy is among the most ancient of sciences, some of the tools used by modern astronomers are actually many centuries old. One such tool is the **magnitude scale,** which astronomers frequently use to denote the brightness of stars. This scale was introduced in the second century B.C. by the Greek astronomer Hipparchus, who called the brightest stars first-magnitude stars. Stars about half as bright as first-magnitude stars were called second-magnitude stars, and so forth, down to sixth-magnitude stars, the dimmest ones he could see. After telescopes came into use, astronomers extended Hipparchus's magnitude scale to include even dimmer stars.

### Apparent Magnitudes

The magnitudes in Hipparchus's scale are properly called **apparent magnitudes,** because they describe how bright an object *appears* to an Earth-based observer. Apparent magnitude is directly related to apparent brightness.

**CAUTION!** The magnitude scale can be confusing because it works "backward." Keep in mind that the *greater* the apparent magnitude, the *dimmer* the star. A star of apparent magnitude +3 (a third-magnitude star) is dimmer than a star of apparent magnitude +2 (a second-magnitude star).

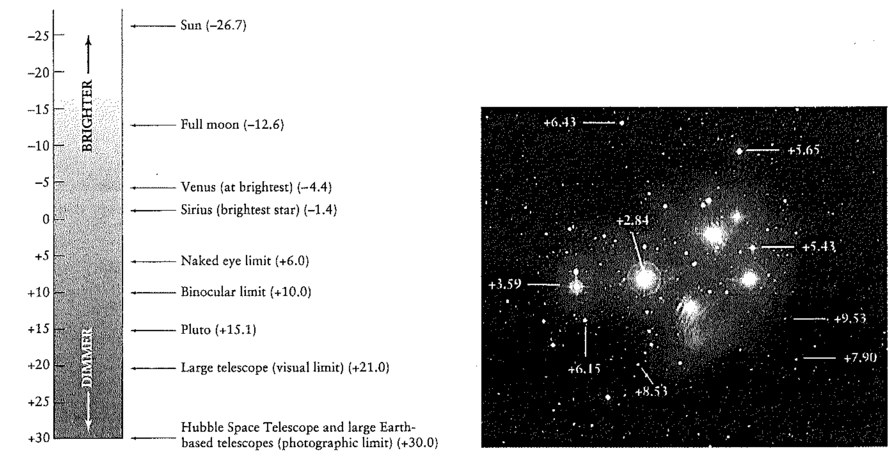


In the nineteenth century, astronomers developed better techniques for measuring the light energy arriving from a star. These measurements showed that a first-magnitude star is about 100 times brighter than a sixth-magnitude star. In other words, it would take 100 stars of magnitude +6 to provide as much light energy as we receive from a single star of magnitude +1. To make computations easier, the magnitude scale was redefined so that a magnitude difference of 5 corresponds exactly to a factor of 100 in brightness. A magnitude difference of 1 corresponds to a factor of 2.512 in brightness, because

$$2.512 \times 2.512 \times 2.512 \times 2.512 \times 2.512 = (2.512)^5 = 100$$

Thus, it takes 2.512 third-magnitude stars to provide as much light as we receive from a single second-magnitude star.

Figure 17-6 illustrates the modern apparent magnitude scale. The dimmest stars visible through a pair of binoculars have an apparent magnitude of +10, and the dimmest stars that can be photographed in a one-hour exposure with the Keck telescopes (see Section 6-2) or the Hubble Space Telescope have apparent magnitude +30. Modern astronomers also use negative numbers to extend Hipparchus's scale to include very bright objects. For example, Sirius, the brightest star in the sky, has an apparent magnitude of −1.43. The Sun, the brightest object in the sky, has an apparent magnitude of −26.7.




(a) Some apparent magnitudes

(b) Apparent magnitudes of stars in the Pleiades

**Figure 17-6**

R I V U X G

The Apparent Magnitude Scale **(a)** Astronomers denote the apparent brightness of objects in the sky by their apparent magnitudes. The greater the apparent magnitude, the dimmer the object. **(b)** This photograph of the Pleiades cluster, located about 120 pc away in the constellation Taurus, shows the apparent magnitudes of some of its stars. Most are too faint to be seen by the naked eye. (David Malin/Anglo-Australian Observatory)

### Absolute Magnitudes

**Apparent magnitude measures a star's brightness; absolute magnitude measures its luminosity**

Apparent magnitude is a measure of a star's apparent brightness as seen from Earth. A related quantity that measures a star's true energy output—that is, its luminosity—is called **absolute magnitude.** This is the apparent magnitude a star would have if it were located exactly 10 parsecs from Earth.

**ANALOGY** If you wanted to compare the light output of two different lightbulbs, you would naturally place them side by side so that both bulbs were the same distance from you. In the absolute magnitude scale, we imagine doing the same thing with stars to compare their luminosities.

If the Sun were moved to a distance of 10 parsecs from Earth, it would have an apparent magnitude of +4.8. The absolute magnitude of the Sun is thus +4.8. The absolute magnitudes of the stars range from approximately +15 for the least luminous to −10 for the most luminous. (*Note:* Like apparent magnitudes, absolute magnitudes work "backward": The *greater* the absolute magnitude, the *less luminous* the star.) The Sun's absolute magnitude is about in the middle of this range.

We saw in Section 17-2 that we can calculate the luminosity of a star if we know its distance and apparent brightness. There is a mathematical relationship between absolute magnitude and luminosity, which astronomers use to convert one to the other as they see fit. It is also possible to rewrite the inverse-square law, which we introduced in Section 17-2, as a mathematical relationship that allows you to calculate a star's absolute magnitude (a measure of its luminosity) from its distance and apparent magnitude (a measure of its apparent brightness). Box 17-3 describes these relationships and how to use them.

Because the "backward" magnitude scales can be confusing, we will use them only occasionally in this book. We will usually speak of a star's luminosity rather than its absolute magnitude and will describe a star's appearance in terms of apparent brightness rather than apparent magnitude. But if you go on to study more about astronomy, you will undoubtedly make frequent use of apparent magnitude and absolute magnitude.

## 17-4 A star's color depends on its surface temperature

WEB LINK 17.7

The image that opens this chapter shows that stars come in different colors. You can see these colors even with the naked eye. For example, you can easily see the red color of Betelgeuse, (the star in the "armpit" of the constellation Orion), and the blue tint of Bellatrix at Orion's other "shoulder" (see Figure 2-2). Colors are most evident for the brightest stars, because human color vision works poorly at low light levels.

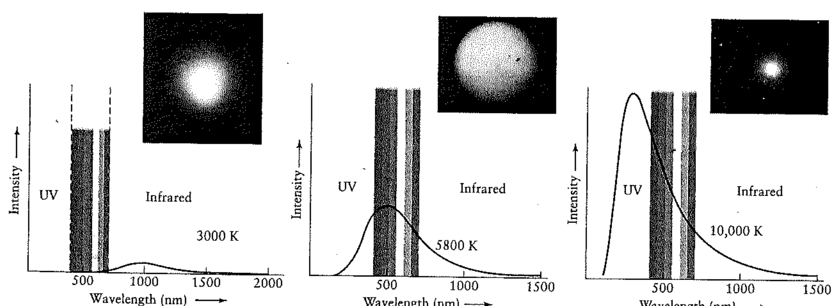
**CAUTION!** It's true that the light from a star will appear redshifted if the star is moving away from you and blueshifted if it's moving toward you. But for even the fastest stars, these color shifts are so tiny that it takes sensitive instruments to measure them. The red color of Betelgeuse and the blue color of Bellatrix are not due to their motions; they are the actual colors of the stars.

### Color and Temperature

We saw in Section 5-3 that a star's color is directly related to its surface temperature. The intensity of light from a relatively cool star peaks at long wavelengths, making the star look red (Figure 17-7*a*). A hot star's intensity curve peaks at shorter wavelengths, so the star looks blue (Figure 17-7*c*). For a star with an intermediate temperature, such as the Sun, the intensity peak is near the middle of the visible spectrum. The human visual system interprets an object with this spectrum of wavelengths as yellowish in color (Figure 17-7*b*). This leads to an important general rule about star colors and surface temperatures:

***Red stars are relatively cold, with low surface temperatures; blue stars are relatively hot, with high surface temperatures.***

Figure 17-7 shows that astronomers can accurately determine the surface temperature of a star by carefully measuring its color.




**(a) A cool star with surface temperature 3000 K emits much more red light than blue light, and so appears red.**

**(b) A warmer star with surface temperature 5800 K (like the Sun) emits similar amounts of all visible wavelengths, and so appears yellow-white.**

**(c) A hot star with surface temperature 10,000 K emits much more blue light than red light, and so appears blue.**

**Figure 17-7**

Temperature and Color These graphs show the intensity of light emitted by three hypothetical stars plotted against wavelength (compare with Figure 5-11). The rainbow band indicates the range of visible wavelengths. The star's apparent color depends on whether the intensity curve has larger values at the short-wavelength or long-wavelength end of the visible spectrum. The insets show stars of about these surface temperatures. UV stands for ultraviolet, which extends from 10 to 400 nm. See Figure 3-4 for more on wavelengths of the spectrum. (Inset a: Andrea Dupree/Harvard-Smithsonian CFA, Ronald Gilliland/STScI, NASA and ESA; inset b: NSO/AURA/NSF; inset c: Till Credner, Allthesky.com)



BOX 17-3 Tools of the Astronomer's Trade

## Apparent Magnitude and Absolute Magnitude

Astronomers commonly express a star's apparent brightness in terms of apparent magnitude (denoted by a lowercase $m$), and the star's luminosity in terms of absolute magnitude (denoted by a capital $M$). While we do not use these quantities extensively in this book, it is useful to know a few simple relationships involving them.

Consider two stars, labeled 1 and 2, with apparent magnitudes $m_1$ and $m_2$ and brightnesses $b_1$ and $b_2$, respectively. The *ratio* of their apparent brightnesses ($b_1/b_2$) corresponds to a *difference* in their apparent magnitudes ($m_2 - m_1$). As we learned in Section 17-3, each step in magnitude corresponds to a factor of 2.512 in brightness; we receive 2.512 times more energy per square meter per second from a third-magnitude star than from a fourth-magnitude star. This idea was used to construct the following table:

| Apparent magnitude difference ($m_2 - m_1$) | Ratio of apparent brightness ($b_1/b_2$) |
|---|---|
| 1 | 2.512 |
| 2 | $(2.512)^2 = 6.31$ |
| 3 | $(2.512)^3 = 15.85$ |
| 4 | $(2.512)^4 = 39.82$ |
| 5 | $(2.512)^5 = 100$ |
| 10 | $(2.512)^{10} = 10^4$ |
| 15 | $(2.512)^{15} = 10^6$ |
| 20 | $(2.512)^{20} = 10^8$ |

A simple equation relates the difference between two stars' apparent magnitudes to the ratio of their brightnesses:

**Magnitude difference related to brightness ratio**

$$m_2 - m_1 = 2.5 \log\left(\frac{b_1}{b_2}\right)$$

$m_1$, $m_2$ = apparent magnitudes of stars 1 and 2

$b_1$, $b_2$ = apparent brightnesses of stars 1 and 2

In this equation, $\log(b_1/b_2)$ is the logarithm of the brightness ratio. The logarithm of $1000 = 10^3$ is 3, the logarithm of $10 = 10^1$ is 1, and the logarithm of $1 = 10^0$ is 0.

EXAMPLE: At their most brilliant, Venus has a magnitude of about −4 and Mercury has a magnitude of about −2. How many times brighter are these planets than the dimmest stars visible to the naked eye, with a magnitude of +6?

**Situation:** In each case we want to find a ratio of two apparent brightnesses (the brightness of Venus or Mercury compared to that of the dimmest naked-eye stars).

**Tools:** In each case we will convert a *difference* in apparent magnitude between the planet and the naked-eye star into a *ratio* of their brightnesses.

**Answer:** The magnitude difference between Venus and the dimmest stars visible to the naked eye is $+6 - (-4) = 10$. From the table, this difference corresponds to a brightness ratio of $(2.512)^{10} = 10^4 = 10{,}000$, so Venus at it most brilliant is 10,000 times brighter than the dimmest naked-eye stars.

The magnitude difference between Mercury and the dimmest naked-eye stars is $+4 - (-4) = 8$. While this value is not in the table, you can see that the corresponding ratio of brightnesses is $(2.512)^8 = (2.512)^{5+3} = (2.512)^5 \times (2.512)^3$. From the table, $(2.512)^5 = 100$ and $(2.512)^3 = 15.85$, so the ratio of brightnesses is $100 \times 15.85 = 1585$. Hence, Mercury at its most brilliant is 1585 times brighter than the dimmest stars visible to the naked eye.

**Review:** Can you show that when at their most brilliant, Venus is 6.31 times brighter than Mercury? (*Hint:* No multiplication or division is required—just notice the difference in apparent magnitude between Venus and Mercury, and consult the table.)

EXAMPLE: The variable star RR Lyrae in the constellation Lyra (the Harp) periodically doubles its light output. By how much does its apparent magnitude change?

**Situation:** We are given a ratio of two brightnesses (the star at its maximum is twice as bright as at its minimum). Our goal is to find the corresponding difference in apparent magnitude.

**Tools:** We let 1 denote the star at its maximum brightness and 2 denote the same star at its dimmest, so the ratio of brightnesses is $b_1/b_2 = 2$. We then use the equation $m_2 - m_1 = 2.5 \log(b_1/b_2)$ to solve for the apparent magnitude difference $m_2 - m_1$.

**Answer:** Using a calculator, we find $m_2 - m_1 = 2.5 \log(2) = 2.5 \times 0.30 = 0.75$. RR Lyrae therefore varies periodically in brightness by 0.75 magnitude.

**Review:** Our answer means that at its dimmest, RR Lyrae has an apparent magnitude $m_2$ that is 0.75 *greater* than its apparent magnitude $m_1$ when it is brightest. (Remember that a greater value of apparent magnitude means the star is dimmer, not brighter!)

*(continued on the next page)*

## BOX 17-3 (continued)

The inverse-square law relating a star's apparent brightness and luminosity can be rewritten in terms of the star's apparent magnitude ($m$), absolute magnitude ($M$), and distance from Earth ($d$). This can be expressed as an equation:

**Relation between a star's apparent magnitude and absolute magnitude**

$$m - M = 5 \log d - 5$$

$m$ = star's apparent magnitude

$M$ = star's absolute magnitude

$d$ = distance from Earth to the star in parsecs

In this expression $m - M$ is called the **distance modulus**, and $\log d$ means the logarithm of the distance $d$ in parsecs. For convenience, the following table gives the values of the distance $d$ corresponding to different values of $m - M$.

| Distance modulus $m - M$ | Distance $d$ (pc) |
|---|---|
| −4 | 1.6 |
| −3 | 2.5 |
| −2 | 4.0 |
| −1 | 6.3 |
| 0 | 10 |
| 1 | 16 |
| 2 | 25 |
| 3 | 40 |
| 4 | 63 |
| 5 | 100 |
| 10 | $10^3$ |
| 15 | $10^4$ |
| 20 | $10^5$ |

This table shows that if a star is less than 10 pc away, its distance modulus $m - M$ is negative. That is, its apparent magnitude ($m$) is less than its absolute magnitude ($M$). If the star is more than 10 pc away, $m - M$ is positive and $m$ is greater than $M$. As an example, the star ε (epsilon) Indi, which is in the direction of the southern constellation Indus, has apparent magnitude $m = +4.7$. It is 3.6 pc away, which is less than 10 pc, so its apparent magnitude is less than its absolute magnitude.

EXAMPLE: Find the absolute magnitude of ε Indi.

**Situation:** We are given the distance to ε Indi ($d = 3.6$ pc) and its apparent magnitude ($m = +4.7$). Our goal is to find the star's absolute magnitude $M$.

**Tools:** We use the formula $m - M = 5 \log d - 5$ to solve for $M$.

**Answer:** Since $d = 3.6$ pc, we use a calculator to find $\log d = \log 3.6 = 0.56$. Therefore the star's distance modulus is $m - M = 5(0.56) - 5 = -2.2$, and the star's absolute magnitude is $M = m - (-2.2) = +4.7 + 2.2 = +6.9$.

**Review:** As a check on our calculations, note that this star's distance modulus $m - M = -2.2$ is less than zero, as it should be for a star less than 10 pc away. Note that our Sun has absolute magnitude +4.8; ε Indi has a greater absolute magnitude, so it is less luminous than the Sun.

EXAMPLE: Suppose you were viewing the Sun from a planet orbiting another star 100 pc away. Could you see it without using a telescope?

**Situation:** We learned in the preceding examples that the Sun has absolute magnitude $M = +4.8$ and that the dimmest stars visible to the naked eye have apparent magnitude $m = +6$. Our goal is to determine whether the Sun would be visible to the naked eye at a distance of 100 pc.

**Tools:** We use the relationship $m - M = 5 \log d - 5$ to find the Sun's apparent magnitude at $d = 100$ pc. If this is greater than +6, the Sun would not be visible at that distance. (Remember that the greater the apparent magnitude, the dimmer the star.)

**Answer:** From the table, at $d = 100$ pc the distance modulus is $m - M = 5$. So, as seen from this distant planet, the Sun's apparent magnitude would be $m = M + 5 = +4.8 + 5 = +9.8$. This is greater than the naked-eye limit $m = +6$, so the Sun could not be seen.

**Review:** The Sun is by far the brightest object in Earth's sky. But our result tells us that to an inhabitant of a planetary system 100 pc away—a rather small distance in a galaxy that is thousands of parsecs across—our own Sun would be just another insignificant star, visible only through binoculars or a telescope.

The magnitude system is also used by astronomers to express the colors of stars as seen through different filters, as we describe in Section 17-4. For example, rather than quantifying a star's color by the color ratio $b_V/b_B$ (a star's apparent brightness as seen through a V filter divided by the brightness through a B filter), astronomers commonly use the *color index* B–V, which is the difference in the star's apparent magnitude as measured with these two filters. We will not use this system in this book, however (but see Advanced Questions 53 and 54).

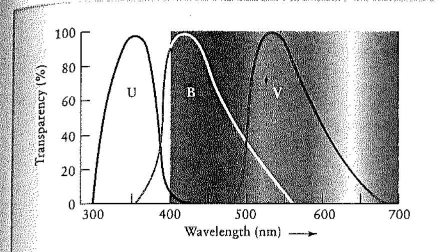



**Figure 17-8**

**U, B, and V Filters** This graph shows the wavelengths to which the standard filters are transparent. The U filter is transparent to near-ultraviolet light. The B filter is transparent to violet, blue, and green light, while the V filter is transparent to green and yellow light. By measuring the apparent brightness of a star with each of these filters and comparing the results, an astronomer can determine the star's surface temperature.

To measure color, the star's light is collected by a telescope and passed through various color-filters. For example, a red filter passes red light while blocking other wavelengths. The filtered light is then collected by a light-sensitive device such as a CCD (see Section 6-4). The process is then repeated with each of the filters in the set. The star's image will have a different brightness through each colored filter, and by comparing these brightnesses astronomers can find the wavelength at which the star's intensity curve has its peak—and hence the star's temperature.

**Astronomers use a set of filters in their telescopes to measure the surface temperatures of stars**

### UBV Photometry

Let's look at this procedure in more detail. The most commonly used filters are called U, B, and V, and the technique that uses them is called **UBV photometry.** Each filter is transparent to a different band of wavelengths: the ultraviolet (U), the blue (B), and the yellow-green (V, for visual) region in and around the visible spectrum (Figure 17-8). The transparency of the V filter mimics the sensitivity of the human eye.

To determine a star's temperature using UBV photometry, the astronomer first measures the star's brightness through each of the filters individually. This gives three apparent brightnesses for the star, designated $b_U$, $b_B$, and $b_V$. The astronomer then compares the intensity of starlight in neighboring wavelength bands by taking the ratios of these brightnesses: $b_V/b_B$ and $b_B/b_U$. Table 17-1 gives values for these **color ratios** for several stars with different surface temperatures.

If a star is very hot, its radiation is skewed toward short, ultraviolet wavelengths as in Figure 17-7*c*. This makes the star dim through the V filter, brighter through the B filter, and brightest through the U filter. Hence, for a hot star $b_V$ is less than $b_B$, which in turn is less than $b_U$, and the ratios $b_V/b_B$ and $b_B/b_U$ are both less than 1. One such star is Bellatrix (see Table 17-1), which has a surface temperature of 21,500 K.

In contrast, if a star is cool, its radiation peaks at long wavelengths as in Figure 17-7*a*. Such a star appears brightest through the V filter, dimmer through the B filter, and dimmest through the U filter (see Figure 17-8). In other words, for a cool star $b_V$ is greater than $b_B$, which in turn is greater than $b_U$. Hence, the ratios $b_V/b_B$ and $b_B/b_U$ will both be greater than 1. The star Betelgeuse (surface temperature 3500 K) is an example.

You can see these differences between hot and cool stars in parts *a* and *c* of Figure 6-27, which show the constellation Orion at ultraviolet wavelengths (a bit shorter than those transmitted by the U filter) and at visible wavelengths that approximate the transmission of a V filter. The hot star Bellatrix is brighter in the ultraviolet image (Figure 6-27*a*) than at visible wavelengths (Figure 6-27*c*). (Figure 6-27*d* shows the names of the stars.) The situation is reversed for the cool star Betelgeuse: It is bright at visible wavelengths, but at ultraviolet wavelengths it is too dim to show up in the image.

Figure 17-9 graphs the relationship between a star's $b_V/b_B$ color ratio and its temperature. If you know the value of the $b_V/b_B$ color ratio for a given star, you can use this graph to find the star's surface temperature. As an example, for the Sun $b_V/b_B$

**Table 17-1** Colors of Selected Stars

| Star | Surface temperature (K) | $b_V/b_B$ | $b_B/b_U$ | Apparent color |
|---|---|---|---|---|
| Bellatrix (γ Orionis) | 21,500 | 0.81 | 0.45 | Blue |
| Regulus (α Leonis) | 12,000 | 0.90 | 0.72 | Blue-white |
| Sirius (α Canis Majoris) | 9400 | 1.00 | 0.96 | Blue-white |
| Megrez (δ Ursae Majoris) | 8630 | 1.07 | 1.07 | White |
| Altair (α Aquilae) | 7800 | 1.23 | 1.08 | Yellow-white |
| Sun | 5800 | 1.87 | 1.17 | Yellow-white |
| Aldebaran (α Tauri) | 4000 | 4.12 | 5.76 | Orange |
| Betelgeuse (α Orionis) | 3500 | 5.55 | 6.66 | Red |

*Source: J.-C. Mermilliod, B. Hauck, and M. Mermilliod, University of Lausanne*

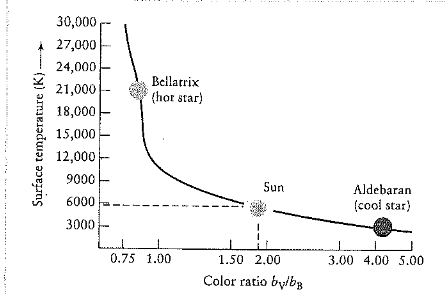



**Figure 17-9**

Temperature, Color, and Color Ratio The $b_V/b_B$ color ratio is the ratio of a star's apparent brightnesses through a V filter and through a B filter. This ratio is small for hot, blue stars but large for cool, red stars. After measuring a star's brightness with the B and V filters, an astronomer can estimate the star's surface temperature from a graph like this one.

equals 1.87, which corresponds to a surface temperature of 5800 K.

**CAUTION!** As we will see in Chapter 18, tiny dust particles that pervade interstellar space cause distant stars to appear redder than they really are. (In the same way, particles in Earth's atmosphere make the setting Sun look redder; see Box 5-4.) Astronomers must take this reddening into account whenever they attempt to determine a star's surface temperature from its color ratios. A star's spectrum provides a more precise measure of a star's surface temperature, as we will see next. But it is quicker and easier to observe a star's colors with a set of U, B, and V filters than it is to take the star's spectrum with a spectrograph.

## 17-5 The spectra of stars reveal their chemical compositions as well as their surface temperatures

We have seen how the color of a star's light helps astronomers determine its surface temperature. To determine the other properties of a star, astronomers must analyze the spectrum of its light in more detail. This technique of *stellar spectroscopy* began in 1817 when Joseph Fraunhofer, a German instrument maker, attached a spectroscope to a telescope and pointed it toward the stars. Fraunhofer had earlier observed that the Sun has an absorption line spectrum—that is, a continuous spectrum with dark absorption lines (see Section 5-6). He found that stars have the same kind of spectra, which reinforces the idea that our Sun is a rather typical star. But Fraunhofer also found that the pattern of absorption lines is different for different stars.

### Classifying Stars: Absorption Line Spectra and Spectral Classes

We see an absorption line spectrum when a cool gas lies between us and a hot, glowing object (recall Figure 5-16). The light from the hot, glowing object itself has a continuous spectrum. In the case of a star, light with a continuous spectrum is produced at low-lying levels of the star's atmosphere where the gases are hot and dense. The absorption lines are created when this light flows outward through the upper layers of the star's atmosphere. Atoms in these cooler, less dense layers absorb radiation at specific wavelengths, which depend on the specific kinds of atoms present—hydrogen, helium, or other elements—and on whether or not the atoms are ionized. Absorption lines in the Sun's spectrum are produced in this same way (see Section 16-5).

WEB LINK 17.8 Some stars have spectra in which the Balmer absorption lines of hydrogen are prominent. But in the spectra of other stars, including the Sun, the Balmer lines are nearly absent and the dominant absorption lines are those of heavier elements such as calcium, iron, and sodium. Still other stellar spectra are dominated by broad absorption lines caused by molecules, such as titanium oxide, rather than single atoms. To cope with this diversity, astronomers group similar-appearing stellar spectra into **spectral classes**. In a popular classification scheme that emerged in the late 1890s, a star was assigned a letter from A through O according to the strength or weakness of the hydrogen Balmer lines in the star's spectrum.

Nineteenth-century science could not explain why or how the spectral lines of a particular chemical are affected by the temperature and density of the gas. Nevertheless, a team of astronomers at the Harvard College Observatory forged ahead with a monumental project of examining the spectra of hundreds of thousands of stars. Their goal was to develop a system of spectral classification in which all spectral features, not just Balmer lines, change smoothly from one spectral class to the next.

**Deciphering the information in starlight took the painstaking work of generations of astronomers**

The Harvard project was financed by the estate of Henry Draper, a wealthy New York physician and amateur astronomer who in 1872 became the first person to photograph stellar absorption lines. Researchers on the project included Edward C. Pickering, Williamina Fleming, Antonia Maury, and Annie Jump Cannon (Figure 17-10). As a result of their efforts, many of the original A-through-O classes were dropped and others were consolidated. The remaining spectral classes were reordered in the sequence **OBAFGKM**. You can remember this sequence with the mnemonic: "Oh, *Be A* Fine Girl (or Guy), *Kiss Me*!"

### Refining the Classification: Spectral Types

Cannon refined the original OBAFGKM sequence into smaller steps called **spectral types**. These steps are indicated by attaching an integer from 0 through 9 to the original letter. For example, the spectral class F includes spectral types F0, F1, F2, . . . , F8, F9, which are followed by the spectral types G0, G1, G2, . . . , G8, G9, and so on.

**Figure 17-10** R I V U X G

WEB LINK 17.9 Classifying the Spectra of the Stars The modern scheme of classifying stars by their spectra was developed at the Harvard College Observatory in the late nineteenth century. A team of women astronomers led by Edward C. Pickering and Williamina Fleming (standing) analyzed hundreds of thousands of spectra. Social conventions of the time prevented most women astronomers from using research telescopes or receiving salaries comparable to men's. (Harvard College Observatory)

Figure 17-11 shows representative spectra of several spectral types. The strengths of spectral lines change gradually from one spectral type to the next. For example, the Balmer absorption lines of hydrogen become increasingly prominent as you go from spectral type B0 to A0. From A0 onward through the F and G classes, the hydrogen lines weaken and almost fade from view. The Sun, whose spectrum is dominated by calcium and iron, is a G2 star.

The Harvard project culminated in the *Henry Draper Catalogue*, published between 1918 and 1924. It listed 225,300 stars, each of which Cannon had personally classified. Meanwhile, physicists had been making important discoveries about the structure of atoms. Ernest Rutherford had shown that atoms have nuclei (recall Figure 5-19), and Niels Bohr made the remarkable hypothesis that electrons circle atomic nuclei along discrete orbits (see Figure 5-22). These advances gave scientists the conceptual and mathematical tools needed to understand stellar spectra.

In the 1920s, the Harvard astronomer Cecilia Payne and the Indian physicist Meghnad Saha demonstrated that the OBAFGKM spectral sequence is actually a sequence in temperature. The hottest stars are O stars. Their absorption lines can occur only if these stars have surface temperatures above 25,000 K. M stars are the coolest stars. The spectral features of M stars are consistent with stellar surface temperatures of about 3000 K.

### Why Surface Temperature Affects Stellar Spectra

To see why the appearance of a star's spectrum is profoundly affected by the star's surface temperature, consider the Balmer lines of hydrogen. Hydrogen is by far the most abundant element in the universe, accounting for about three-quarters of the mass of a typical star. Yet the Balmer lines do not necessarily show up in a star's spectrum. As we saw in Section 5-8, Balmer absorption lines are produced when an electron in the $n = 2$ orbit of hydrogen is lifted into a higher orbit by absorbing a photon with the right amount of energy (see Figure 5-24). If the star is much hotter than 10,000 K, the photons pouring out of the star's interior have such high energy that they easily knock electrons out of hydrogen atoms in the star's atmosphere. This process ionizes the gas. With its only electron torn away, a hydrogen atom cannot produce absorption lines. Hence, the Balmer lines will be relatively weak in the spectra of such hot stars, such as the hot O and B2 stars in Figure 17-11.

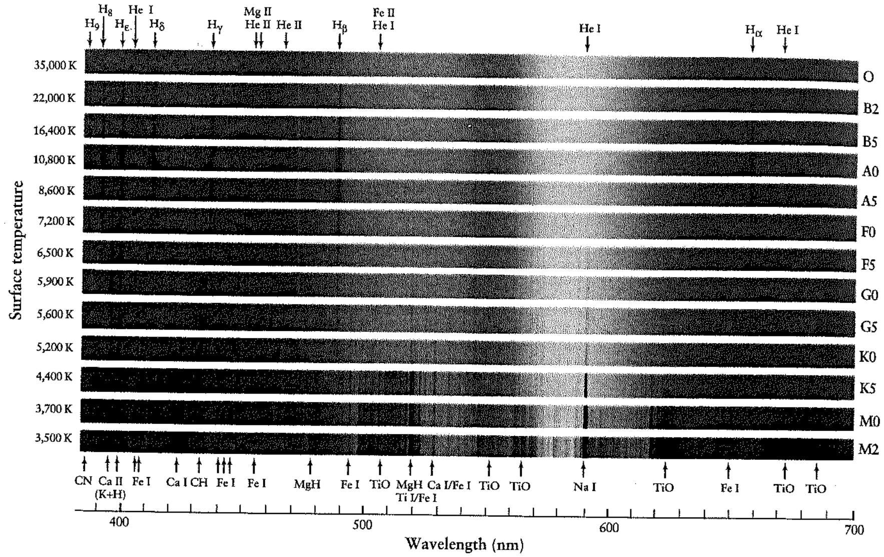
Conversely, if the star's atmosphere is much cooler than 10,000 K, almost all the hydrogen atoms are in the lowest ($n = 1$) energy state. Most of the photons passing through the star's atmosphere possess too little energy to boost electrons up from the $n = 1$ to the $n = 2$ orbit of the hydrogen atoms. Hence, very few of these atoms will have electrons in the $n = 2$ orbit, and only these few can absorb the photons characteristic of the Balmer lines. As a result, these lines are nearly absent from the spectrum of a cool star. (You can see this in the spectra of the cool M0 and M2 stars in Figure 17-11.)

For the Balmer lines to be prominent in a star's spectrum, the star must be hot enough to excite the electrons out of the ground state but not so hot that all the hydrogen atoms become ionized. A stellar surface temperature of about 9000 K produces the strongest hydrogen lines; this is the case for the stars of spectral types A0 and A5 in Figure 17-11.

LOOKING DEEPER 17.1 Every other type of atom or molecule also has a characteristic temperature range in which it produces prominent absorption lines in the observable part of the spectrum. Figure 17-12 shows the relative strengths of absorption lines produced by different chemicals. By measuring the details of these lines in a given star's spectrum, astronomers can accurately determine that star's surface temperature.

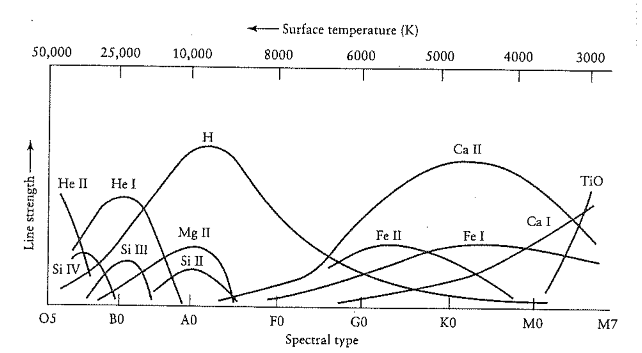
For example, the spectral lines of neutral (that is, un-ionized) helium are strong around 25,000 K. At this temperature, photons have enough energy to excite helium atoms without tearing away the electrons altogether. In stars hotter than about 30,000 K, helium atoms become singly ionized, that is, they lose one of their two electrons. The remaining electron produces a set of spectral lines that is recognizably different from those of neutral helium. Hence, when the spectral lines of singly ionized helium appear in a star's spectrum, we know that the star's surface temperature is greater than 30,000 K.

Astronomers use the term **metals** to refer to all elements other than hydrogen and helium. This idiosyncratic use of the term "metal" is quite different from the definition used by chemists and other scientists. To a chemist, sodium and iron are metals but carbon and oxygen are not; to an astronomer, all of these substances are metals. In this terminology, metals dominate the




## Figure 17-11 R I V U X G

Principal Types of Stellar Spectra Stars of different spectral classes and different surface temperatures have spectra dominated by different absorption lines. Notice how the Balmer lines of hydrogen ($H_\alpha$, $H_\beta$, $H_\gamma$, and $H_\delta$) are strongest for hot stars of spectral class A, while absorption lines due to calcium (Ca) are strongest in cooler K and M stars. The spectra of M stars also have broad, dark bands caused by molecules of titanium oxide (TiO), which can only exist at relatively low temperatures. A roman numeral after a chemical symbol shows whether the absorption line is caused by un-ionized atoms (roman numeral I) or by atoms that have lost one electron (roman numeral II). (R. Bell, University of Maryland, and M. Briley, University of Wisconsin at Oshkosh)




## Figure 17-12

The Strengths of Absorption Lines Each curve in this graph peaks at the stellar surface temperature for which that chemical's absorption line is strongest. For example, hydrogen (H) absorption lines are strongest in A stars with surface temperatures near 10,000 K. Roman numeral I denotes neutral, un-ionized atoms; II, III, and IV denote atoms that are singly, doubly, or triply ionized (that is, have lost one, two, or three electrons).



spectra of stars cooler than 10,000 K. Ionized metals are prominent for surface temperatures between 6000 and 8000 K, while neutral metals are strongest between approximately 5500 and 4000 K.

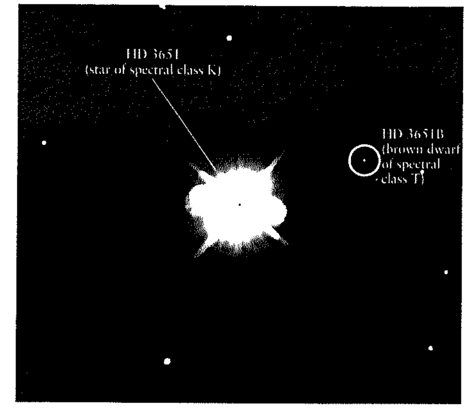
Below 4000 K, certain atoms in a star's atmosphere combine to form molecules. (At higher temperatures atoms move so fast that when they collide, they bounce off each other rather than "sticking together" to form molecules.) As these molecules vibrate and rotate, they produce bands of spectral lines that dominate the star's spectrum. Most noticeable are the lines of titanium oxide (TiO), which are strongest for surface temperatures of about 3000 K.

### Spectral Classes for Brown Dwarfs

WEBLINK 17.10

Since 1995 astronomers have found a number of stars with surface temperatures even lower than those of spectral class M. Strictly speaking, these are not stars but **brown dwarfs**, which we introduced in Section 8-6. Brown dwarfs are too small to sustain thermonuclear fusion in their cores. Instead, these "substars" glow primarily from the heat released by Kelvin-Helmholtz contraction, which we described in Section 16-1. (They do undergo fusion reactions for a brief period during their evolution.) Brown dwarfs are so cold that they are best observed with infrared telescopes (see Figure 17-13). Such observations reveal that brown dwarf spectra have a rich variety of absorption lines due to molecules. Some of these molecules actually form into solid grains in a brown dwarf's atmosphere.

To describe brown dwarf spectra, astronomers have defined two new spectral classes, L and T. Thus, the modern spectral sequence of stars and brown dwarfs from hottest to coldest surface temperature is OBAFGKMLT. (Can you think of a new mnemonic that includes L and T?) For example, Figure 17-13 shows a star of spectral class K and a brown dwarf of spectral class T. Table 17-2 summarizes the relationship between the temperature and spectra of stars and brown dwarfs.




## Figure 17-13 R I V U X G

An Infrared Image of Brown Dwarf HD 3651B The star HD 3561 is of spectral class K, with a surface temperature of about 5200 K. ("HD" refers to the *Henry Draper Catalogue*.) In 2006 it was discovered that HD 3651 is orbited by a brown dwarf named HD 3651B with a surface temperature between 800 and 900 K and a luminosity just 1/300,000 that of the Sun. The brown dwarf emits most of its light at infrared wavelengths, so an infrared telescope was used to record this image. The hotter and more luminous star HD 3651 is greatly overexposed in this image and appears much larger than its true size. HD 3651 and HD 3651B are both 11 pc (36 ly) from Earth in the constellation Pisces (the Fish); the other stars in this image are much farther away. (M. Mugrauer and R. Neuhäuser, U. of Jena; A. Seifahrt, ESO; and T. Mazeh, Tel Aviv U.)

**Table 17-2 The Spectral Sequence**

| Spectral class | Color | Temperature (K) | Spectral lines | Examples |
|---|---|---|---|---|
| O | Blue-violet | 30,000–50,000 | Ionized atoms, especially helium | Naos (ζ Puppis), Mintaka (δ Orionis) |
| B | Blue-white | 11,000–30,000 | Neutral helium, some hydrogen | Spica (α Virginis), Rigel (β Orionis) |
| A | White | 7500–11,000 | Strong hydrogen, some ionized metals | Sirius (α Canis Majoris), Vega (α Lyrae) |
| F | Yellow-white | 5900–7500 | Hydrogen and ionized metals such as calcium and iron | Canopus (α Carinae), Procyon (α Canis Minoris) |
| G | Yellow | 5200–5900 | Both neutral and ionized metals, especially ionized calcium | Sun, Capella (α Aurigae) |
| K | Orange | 3900–5200 | Neutral metals | Arcturus (α Boötis), Aldebaran (α Tauri) |
| M | Red-orange | 2500–3900 | Strong titanium oxide and some neutral calcium | Antares (α Scorpii), Betelgeuse (α Orionis) |
| L | Red | 1300–2500 | Neutral potassium, rubidium, and cesium, and metal hydrides | Brown dwarf Teide 1 |
| T | Red | below 1300 | Strong neutral potassium and some water ($H_2O$) | Brown dwarfs Gliese 229B, HD 3651B |

When the effects of temperature are accounted for, astronomers find that *all* stars have essentially the same chemical composition. We can state the results as a general rule:

***By mass, almost all stars (including the Sun) and brown dwarfs are about three-quarters hydrogen, one-quarter helium, and 1% or less metals.***

Our Sun is about 1% metals by mass, as are most of the stars you can see with the naked eye. But some stars have an even lower percentage of metals. We will see in Chapter 19 that these seemingly minor differences tell an important tale about the life stories of stars.

## 17-6 Stars come in a wide variety of sizes

With even the best telescopes, stars appear as nothing more than bright points of light. On a photograph or CCD image, brighter stars appear larger than dim ones (see Figures 17-3, 17-6*b*, and 17-13), but these apparent sizes give no indication of the star's actual size. To determine the size of a star, astronomers combine information about its luminosity (determined from its distance and apparent brightness) and its surface temperature (determined from its spectral type). In this way, they find that some stars are quite a bit smaller than the Sun, while others are a thousand times larger.

**A star's radius can be calculated if we know its luminosity and surface temperature**

### Calculating the Radii of Stars

The key to finding a star's radius from its luminosity and surface temperature is the Stefan-Boltzmann law (see Section 5-4). This law says that the amount of energy radiated per second from a square meter of a blackbody's surface—that is, the energy flux ($F$)—is proportional to the fourth power of the temperature of that surface ($T$), as given by the equation $F = \sigma T^4$. This equation applies very well to stars, whose spectra are quite similar to that of a perfect blackbody. (Absorption lines, while important for determining the star's chemical composition and surface temperature, make only relatively small modifications to a star's blackbody spectrum.)

A star's luminosity is the amount of energy emitted per second from its entire surface. This quantity equals the energy flux $F$ multiplied by the total number of square meters on the star's surface (that is, the star's surface area). We expect that most stars are nearly spherical, like the Sun, so we can use the formula for the surface area of a sphere. The formula is $4\pi R^2$, where $R$ is the star's radius (the distance from its center to its surface). Multiplying together the formulas for energy flux and surface area, we can write the star's luminosity as follows:

**Relationship between a star's luminosity, radius, and surface temperature**

$$L = 4\pi R^2 \sigma T^4$$

$L$ = star's luminosity, in watts

$R$ = star's radius, in meters

$\sigma$ = Stefan-Boltzmann constant = $5.67 \times 10^{-8}$ W m$^{-2}$ K$^{-4}$

$T$ = star's surface temperature, in kelvins

This equation says that a relatively cool star (low surface temperature $T$), for which the energy flux is quite low, can nonetheless be very luminous if it has a large enough radius $R$. Alternatively, a relatively hot star (large $T$) can have a very low luminosity if the star has only a little surface area (small $R$).

Box 17-4 describes how to use the above equation to calculate a star's radius if its luminosity and surface temperature are known. We can express the idea behind these calculations in terms of the following general rule:

***We can determine the radius of a star from its luminosity and surface temperature. For a given luminosity, the greater the surface temperature, the smaller the radius must be. For a given surface temperature, the greater the luminosity, the larger the radius must be.***

ANALOGY In a similar way, a roaring campfire can emit more light than a welder's torch. The campfire is at a lower temperature than the torch, but has a much larger surface area from which it emits light.

### The Range of Stellar Radii

Using this general rule as shown in Box 17-4, astronomers find that stars come in a wide range of sizes. The smallest stars visible through ordinary telescopes, called *white dwarfs*, are about the same size as Earth. Although their surface temperatures can be very high (25,000 K or more), white dwarfs have so little surface area that their luminosities are very low (less than 0.01 $L_\odot$). The largest stars, called *supergiants*, are a thousand times larger in radius than the Sun and $10^5$ times larger than white dwarfs. If our own Sun were replaced by one of these supergiants, Earth's orbit would lie completely inside the star!

Figure 17-14 summarizes how astronomers determine the distance from Earth, luminosity, surface temperature, chemical

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**BOX 17-4** Tools of the Astronomer's Trade

## Stellar Radii, Luminosities, and Surface Temperatures

Because stars emit light in almost exactly the same fashion as blackbodies, we can use the Stefan-Boltzmann law to relate a star's luminosity ($L$), surface temperature ($T$), and radius ($R$). The relevant equation is

$$L = 4\pi R^2 \sigma T^4$$

As written, this equation involves the Stefan-Boltzmann constant $\sigma$, which is equal to $5.67 \times 10^{-8}$ W m$^{-2}$ K$^{-4}$. In many calculations, it is more convenient to relate everything to the Sun, which is a typical star. Specifically, for the Sun we have $L_\odot = 4\pi R_\odot^2 \sigma T_\odot^4$, where $L_\odot$ is the Sun's luminosity, $R_\odot$ is the Sun's radius, and $T_\odot$ is the Sun's surface temperature (equal to 5800 K). Dividing the general equation for $L$ by this specific equation for the Sun, we obtain

$$\frac{L}{L_\odot} = \left(\frac{R}{R_\odot}\right)^2 \left(\frac{T}{T_\odot}\right)^4$$

This formula is easier to use because the constant $\sigma$ has cancelled out. We can also rearrange terms to arrive at a useful equation for the radius ($R$) of a star:

**Radius of a star related to its luminosity and surface temperature**

$$\frac{R}{R_\odot} = \left(\frac{T_\odot}{T}\right)^2 \sqrt{\frac{L}{L_\odot}}$$

$R/R_\odot$ = ratio of the star's radius to the Sun's radius

$T_\odot/T$ = ratio of the Sun's surface temperature to the star's surface temperature

$L/L_\odot$ = ratio of the star's luminosity to the Sun's luminosity

EXAMPLE: The bright reddish star Betelgeuse in the constellation Orion (see Figure 2-2) is 60,000 times more luminous than the Sun and has a surface temperature of 3500 K. What is its radius?

Situation: We are given the star's luminosity $L = 60{,}000\ L_\odot$ and its surface temperature $T = 3500$ K. Our goal is to find the star's radius $R$.

Tools: We use the above equation to find the ratio of the star's radius to the radius of the Sun, $R/R_\odot$. Note that we also know the Sun's surface temperature, $T_\odot = 5800$ K.

Answer: Substituting these data into the above equation, we get

$$\frac{R}{R_\odot} = \left(\frac{5800\ \text{K}}{3500\ \text{K}}\right)^2 \sqrt{6 \times 10^4} = 670$$

Review: Our result tells us that Betelgeuse's radius is 670 times larger than that of the Sun. The Sun's radius is $6.96 \times 10^5$ km, so we can also express the radius of Betelgeuse as $(670)(6.96 \times 10^5\ \text{km}) = 4.7 \times 10^8$ km, which is more than 3 AU. If Betelgeuse were located at the center of our solar system, it would extend beyond the orbit of Mars!

EXAMPLE: Sirius, the brightest star in the sky, is actually two stars orbiting each other (a binary star). The less luminous star, Sirius B, is a white dwarf that is too dim to see with the naked eye. Its luminosity is 0.0025 $L_\odot$ and its surface temperature is 10,000 K. How large is Sirius B compared to Earth?

Situation: Again we are asked to find a star's radius from its luminosity and surface temperature.

Tools: We use the same equation as in the preceding example.

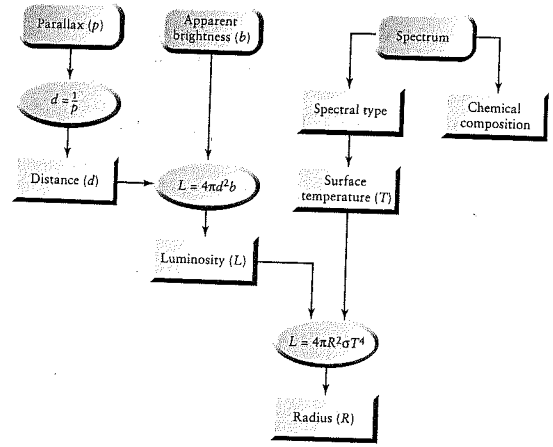
Answer: The ratio of the radius of Sirius B to the Sun's radius is

$$\frac{R}{R_\odot} = \left(\frac{5800\ \text{K}}{10{,}000\ \text{K}}\right)^2 \sqrt{0.0025} = 0.017$$

Since the Sun's radius is $R_\odot = 6.96 \times 10^5$ km, the radius of Sirius B is $(0.017)(6.96 \times 10^5\ \text{km}) = 12{,}000$ km. From Table 7-1, Earth's radius (half its diameter) is 6378 km. Hence, this star is only about twice the radius of Earth.

Review: Sirius B's radius would be large for a terrestrial planet, but it is minuscule for a star. The name *dwarf* is well deserved!

The radii of some stars have been measured with other techniques (see Section 17-11). These other methods yield values consistent with those calculated by the methods we have just described.




INTERACTIVE EXERCISE 17.1

**Figure 17-14**

Finding Key Properties of a Nearby Star This flowchart shows how astronomers determine the properties of a relatively nearby star (one close enough that its parallax can be measured). The rounded purple boxes show the measurements that must be made of the star, the blue ovals show the key equations that are used (from Sections 17-2, 17-5, and 17-6), and the green rectangles show the inferred properties of the stars. A different procedure is followed for more distant stars (see Section 17-8, especially Figure 17-17).

composition, and radius of a star close enough to us so that its parallax can be measured. Remarkably, all of these properties can be deduced from just a few measured quantities: the star's parallax angle, apparent brightness, and spectrum.

## 17-7 Hertzsprung-Russell (H-R) diagrams reveal the different kinds of stars

Astronomers have collected a wealth of data about the stars, but merely having tables of numerical data is not enough. Like all scientists, astronomers want to analyze their data to look for trends and underlying principles. One of the best ways to look for trends in any set of data, whether it comes from astronomy, finance, medicine, or meteorology, is to create a graph showing how one quantity depends on another. For example, investors consult graphs of stock market values versus dates, and weather forecasters make graphs of temperature versus altitude to determine whether thunderstorms are likely to form. Astronomers have found that a particular graph of stellar properties shows that stars fall naturally into just a few categories. This graph, one of the most important in all astronomy, will in later chapters help us understand how stars form, evolve, and eventually die.

### H-R Diagrams

Which properties of stars should we include in a graph? Most stars have about the same chemical composition, but two properties of stars—their luminosities and surface temperatures—differ substantially from one star to another. Stars also come in a wide range of radii, but a star's radius is a secondary property that can be found from the luminosity and surface temperature (as we saw in Section 17-6 and Box 17-4). We also relegate the positions and space velocities of stars to secondary importance. (In a similar way, a physician is more interested in your weight and blood pressure than in where you live or how fast you drive.) We can then ask the following question: What do we learn when we graph the luminosities of stars versus their surface temperatures?

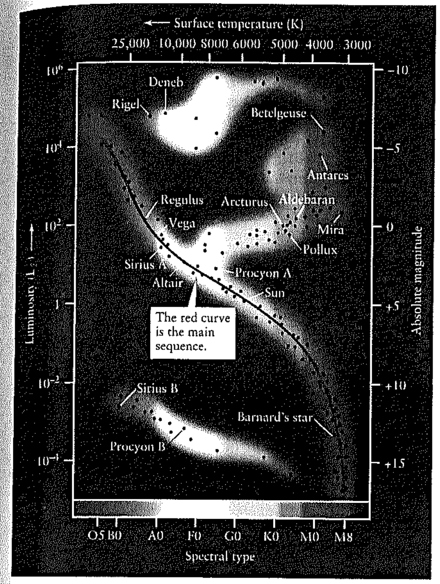
**A graph of luminosity versus surface temperature reveals that stars fall into a few basic categories**

The first answer to this question was given in 1911 by the Danish astronomer Ejnar Hertzsprung. He pointed out that a regular pattern appears when the absolute magnitudes of stars (which measure their luminosities) are plotted against their colors (which measure their surface temperatures). Two years later, the American astronomer Henry Norris Russell independently discovered a similar regularity in a graph using spectral types (another measure of surface temperature) instead of colors. In recognition of their originators, graphs of this kind are today known as **Hertzsprung-Russell diagrams**, or **H-R diagrams** (Figure 17-15).

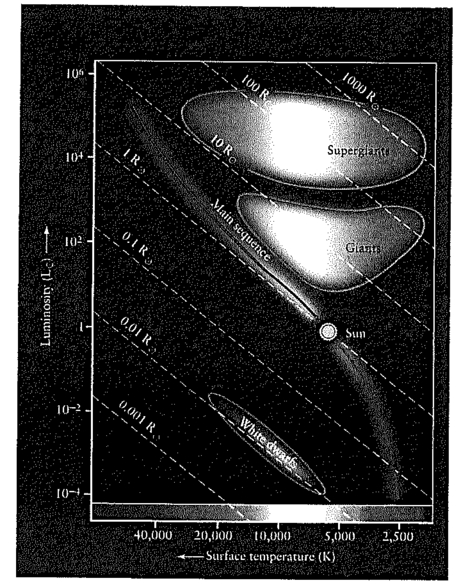
Figure 17-15*a* is a typical Hertzsprung-Russell diagram. Each dot represents a star whose spectral type and luminosity have been determined. The most luminous stars are near the top of the diagram, the least luminous stars near the bottom. Hot stars of spectral classes O and B are toward the left side of the graph and cool stars of spectral class M are toward the right.

**CAUTION!** You are probably accustomed to graphs in which the numbers on the horizontal axis increase as you move to the right. (For example, the business section of a newspaper includes a graph of stock market values versus dates, with later dates to the right of earlier ones.) But on an H-R diagram the temperature scale on the horizontal axis increases toward the *left*. This practice stems from the original diagrams of Hertzsprung and Russell, who placed hot O stars on the left and cool M stars on the right. This arrangement is a tradition that no one has seriously tried to change.






(a) A Hertzsprung-Russell (H-R) diagram




(b) The sizes of stars on an H-R diagram

INTERACTIVE EXERCISE 17.2

**Figure 17-15**

Hertzsprung-Russell (H-R) Diagrams On an H-R diagram, the luminosities (or absolute magnitudes) of stars are plotted against their spectral types (or surface temperatures). **(a)** The data points are grouped in just a few regions on the graph, showing that luminosity and spectral type are correlated. Most stars lie along the red curve called the main sequence. Giants like Arcturus as well as supergiants like Rigel and Betelgeuse are above the main sequence, and white dwarfs like Sirius B are below it. **(b)** The blue curves on this H-R diagram enclose the regions of the diagram in which different types of stars are found. The dashed diagonal lines indicate different stellar radii. For a given stellar radius, as the surface temperature increases (that is, moving from right to left in the diagram), the star glows more intensely and the luminosity increases (that is, moving upward in the diagram). Note that the Sun is intermediate in luminosity, surface temperature, and radius.

### Star Varieties: Main-Sequence Stars, Giants, Supergiants, White Dwarfs, and Brown Dwarfs

The most striking feature of the H-R diagram is that the data points are not scattered randomly over the graph but are grouped in a few distinct regions. The luminosities and surface temperatures of stars do *not* have random values; instead, these two quantities are related!

The band stretching diagonally across the H-R diagram includes about 90% of the stars in the night sky. This band, called **the main sequence,** extends from the hot, luminous, blue stars in the upper left corner of the diagram to the cool, dim, red stars in the lower right corner. A star whose properties place it in this region of an H-R diagram is called a **main-sequence star.** The Sun (spectral type G2, luminosity 1 $L_\odot$, absolute magnitude +4.8) is such a star. We will find that all main-sequence stars are like the Sun in that *hydrogen fusion*—thermonuclear reactions that convert hydrogen into helium (see Section 16-1)—is taking place in their cores.

The upper right side of the H-R diagram shows a second major grouping of data points. Stars represented by these points are both luminous and cool. From the Stefan-Boltzmann law, we know that a cool star radiates much less light per unit of surface area than a hot star. In order for these stars to be as luminous as they are, they must be huge (see Section 17-6), and so they are called **giants**. These stars are around 10 to 100 times larger than the Sun. You can see this size difference in Figure 17-15*b*, which is an H-R diagram to which dashed lines have been added to represent stellar radii. Most giant stars are around 100 to 1000 times more luminous than the Sun and have surface temperatures of about 3000 to 6000 K. Cooler members of this class of stars (those with surface temperatures from about 3000 to 4000 K) are often called **red giants** because they appear reddish. In the image that opens this chapter, the yellowish stars, as well as the red star just left of center, are red giants. A number of red giants can easily be seen with the naked eye, including Aldebaran in the constellation Taurus and Arcturus in Boötes.

A few rare stars are considerably bigger and brighter than typical red giants, with radii up to 1000 $R_{\odot}$. Appropriately enough, these superluminous stars are called **supergiants**. Betelgeuse in Orion (see Box 17-4) and Antares in Scorpius are two supergiants you can find in the nighttime sky. Together, giants and supergiants make up about 1% of the stars in the sky.

Both giants and supergiants have thermonuclear reactions occurring in their interiors, but the character of those reactions and where in the star they occur can be quite different than for a main-sequence star like the Sun. We will study these stars in more detail in Chapters 21 and 22.

The remaining 9% of stars form a distinct grouping of data points toward the lower left corner of the Hertzsprung-Russell diagram. Although these stars are hot, their luminosities are quite low; hence, they must be small. They are appropriately called **white dwarfs**. These stars, which are so dim that they can be seen only with a telescope, are approximately the same size as Earth. As we will learn in Chapter 20, no thermonuclear reactions take place within white dwarf stars. Rather, like embers left from a fire, they are the still-glowing remnants of what were once giant stars.

By contrast, *brown* dwarfs (which lie at the extreme lower right of the main sequence, off the bottom and right-hand edge of Figure 17-15*a* or Figure 17-15*b*) are objects that will never become stars. They are comparable in radius to the planet Jupiter (that is, intermediate in size between Earth and the Sun; see Figure 7-2). The study of brown dwarfs is still in its infancy, but it appears that there may be twice as many brown dwarfs as there are "real" stars.

ANALOGY You can think of white dwarfs as "has-been" stars whose days of glory have passed. In this analogy, a brown dwarf is a "never-will-be."

The existence of fundamentally different types of stars is the first important lesson to come from the H-R diagram. In later chapters we will find that these different types represent various stages in the lives of stars. We will use the H-R diagram as an essential tool for understanding how stars evolve.

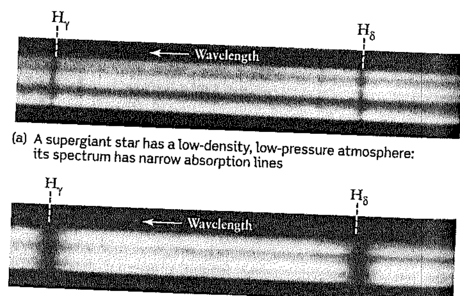
## 17-8 Details of a star's spectrum reveal whether it is a giant, a white dwarf, or a main-sequence star

A star's surface temperature largely determines which lines are prominent in its spectrum. Therefore, classifying stars by spectral type is essentially the same as categorizing them by surface temperature. But as the H-R diagram in Figure 17-15*b* shows, stars of the same surface temperature can have very different luminosities. As an example, a star with surface temperature 5800 K could be a white dwarf, a main-sequence star, a giant, or a supergiant, depending on its luminosity. By examining the details of a star's spectrum, however, astronomers can determine to which of these categories a star belongs. This gives astronomers a tool to determine the distances to stars millions of parsecs away, far beyond the maximum distance that can be measured using stellar parallax.

### Determining a Star's Size from Its Spectrum

Figure 17-16 compares the spectra of two stars of the same spectral type but different luminosity (and hence different size): a B8 supergiant and a B8 main-sequence star. Note that the Balmer lines of hydrogen are narrow in the spectrum of the very large, very luminous supergiant but quite broad in the spectrum of the small, less luminous main-sequence star. In general, for stars of spectral types B through F, the larger and more luminous the star, the narrower its hydrogen lines.

**The smaller a star and the denser its atmosphere, the broader the absorption lines in its spectrum**



(a) A supergiant star has a low-density, low-pressure atmosphere: its spectrum has narrow absorption lines

(b) A main-sequence star has a denser, higher-pressure atmosphere: its spectrum has broad absorption lines

**Figure 17-16** R I V U X G

How a Star's Size Affects Its Spectrum These are the spectra of two stars of the same spectral type (B8) and surface temperature (13,400 K) but different radii and luminosities: **(a)** the B8 supergiant Rigel (luminosity 58,000 $L_{\odot}$) in Orion, and **(b)** the B8 main-sequence star Algol (luminosity 100 $L_{\odot}$) in Perseus. (From W. W. Morgan, P. C. Keenan, and E. Kellman, *An Atlas of Stellar Spectra*)



Fundamentally, these differences between stars of different luminosity are due to differences between the stars' atmospheres, where absorption lines are produced. Hydrogen lines in particular are affected by the density and pressure of the gas in a star's atmosphere. The higher the density and pressure, the more frequently hydrogen atoms collide and interact with other atoms and ions in the atmosphere. These collisions shift the energy levels in the hydrogen atoms and thus broaden the hydrogen spectral lines.

In the atmosphere of a luminous giant star, the density and pressure are quite low because the star's mass is spread over a huge volume. Atoms and ions in the atmosphere are relatively far apart; hence, collisions between them are sufficiently infrequent that hydrogen atoms can produce narrow Balmer lines. A main-sequence star, however, is much more compact than a giant or supergiant. In the denser atmosphere of a main-sequence star, frequent interatomic collisions perturb the energy levels in the hydrogen atoms, thereby producing broader Balmer lines.

### Luminosity Classes

In the 1930s, W. W. Morgan and P. C. Keenan of the Yerkes Observatory of the University of Chicago developed a system of **luminosity classes** based upon the subtle differences in spectral lines. When these luminosity classes are plotted on an H-R diagram (Figure 17-17), they provide a useful subdivision of the star types in the upper right of the diagram. Luminosity classes Ia and Ib are composed of supergiants; luminosity class V includes all the main-sequence stars. The intermediate classes distinguish giant stars of various luminosities. Note that for stars of a given surface temperature (that is, a given spectral type), the *higher* the number of the luminosity class, the *lower* the star's luminosity.

As we will see in Chapters 19 and 20, different luminosity classes represent different stages in the evolution of a star. White dwarfs are not always given a luminosity class of their own; as we mentioned in Section 17-7, they represent a final stage in stellar evolution in which no thermonuclear reactions take place.

Astronomers commonly use a shorthand description that combines a star's spectral type and its luminosity class. For example, the Sun is said to be a G2 V star. The spectral type indicates the star's surface temperature, and the luminosity class indicates its luminosity. Thus, an astronomer knows immediately that any G2 V star is a main-sequence star with a luminosity of about 1 $L_{\odot}$ and a surface temperature of about 5800 K. Similarly, a description of Aldebaran as a K5 III star tells an astronomer that it is a red giant with a luminosity of around 370 $L_{\odot}$ and a surface temperature of about 4000 K.

### Spectroscopic Parallax

A star's spectral type and luminosity class, combined with the information on the H-R diagram, enable astronomers to estimate the star's distance from Earth. As an example, consider the star Pleione in the constellation Taurus. Its spectrum reveals Pleione to be a B8 V star (a hot, blue, main-sequence star, like the one in Figure 17-16*b*). Using Figure 17-17, we can read off that such a star's luminosity is 190 $L_{\odot}$. Given the star's luminosity and its apparent brightness—in the case of Pleione, $3.9 \times 10^{-13}$ of the apparent brightness of the Sun—we can use the inverse-square law to determine its distance from Earth. The mathematical details are worked out in Box 17-2.

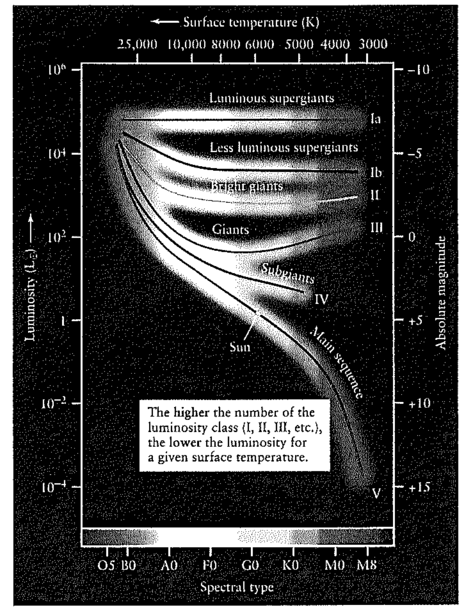



**Figure 17-17**

Luminosity Classes The H-R diagram is divided into regions corresponding to stars of different luminosity classes. (White dwarfs do not have their own luminosity class.) A star's spectrum reveals both its spectral type and its luminosity class; from these, the star's luminosity can be determined.

This method for determining distance, in which the luminosity of a star is found using spectroscopy, is called **spectroscopic parallax**. Figure 17-18 summarizes the method of spectroscopic parallax.

**CAUTION!** The name "spectroscopic parallax" is a bit misleading, because no parallax angle is involved! The idea is that measuring the star's spectrum takes the place of measuring its parallax as a way to find the star's distance. A better name for this method, although not the one used by astronomers, would be "spectroscopic distance determination."

Spectroscopic parallax is an incredibly powerful technique. No matter how remote a star is, this technique allows astronomers

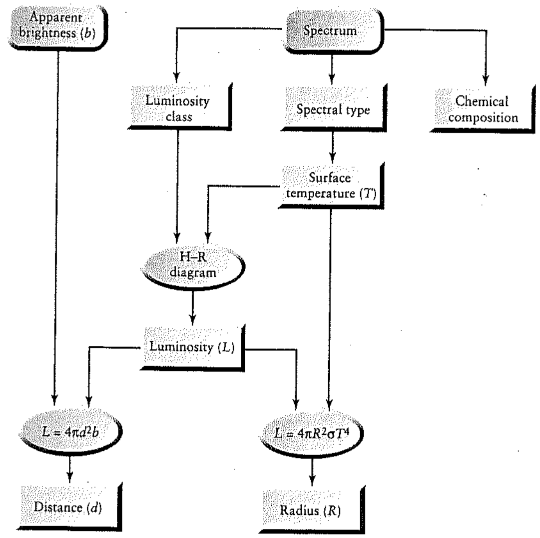



**Figure 17-18**

The Method of Spectroscopic Parallax If a star is too far away, its parallax angle is too small to allow a direct determination of its distance. This flowchart shows how astronomers deduce the properties of such a distant star. Note that the H-R diagram plays a central role in determining the star's luminosity from its spectral type and luminosity class. Just as for nearby stars (see Figure 17-14), the star's chemical composition is determined from its spectrum, and the star's radius is calculated from the luminosity and surface temperature.

to determine its distance, provided only that its spectrum and apparent brightness can be measured. Box 17-2 gives an example of how spectroscopic parallax has been used to find the distance to stars in other galaxies tens of millions of parsecs away. By contrast, we saw in Section 17-1 that "real" stellar parallaxes can be measured only for stars within a few hundred parsecs.

Unfortunately, spectroscopic parallax has its limitations; distances to individual stars determined using this method often have errors greater than 10%. The reason is that the luminosity classes shown in Figure 17-17 are not thin lines on the H-R diagram but are moderately broad bands. Hence, even if a star's spectral type and luminosity class are known, there is still some uncertainty in the luminosity that we read off an H-R diagram. Nonetheless, spectroscopic parallax is often the only means that an astronomer has to estimate the distance to remote stars.

What has been left out of this discussion is *why* different stars have different spectral types and luminosities. One key factor, as we shall see, turns out to be the mass of the star.

## 17-9 Observing binary star systems reveals the masses of stars

TUTORIAL 17.3

We now know something about the sizes, temperatures, and luminosities of stars. To complete our picture of the physical properties of stars, we need to know their masses. In this section, we will see that stars come in a wide range of masses. We will also discover an important relationship between the mass and luminosity of main-sequence stars. This relationship is crucial to understanding why some main-sequence stars are hot and luminous, while others are cool and dim. It will also help us understand what happens to a star as it ages and evolves.

**For main-sequence stars, there is a direct correlation between mass and luminosity**

Determining the masses of stars is not trivial, however. The problem is that there is no practical, direct way to measure the mass of an isolated star. Fortunately for astronomers, about half of the visible stars in the night sky are not isolated individuals. Instead, they are ***multiple-star systems,*** in which two or more stars orbit each other. By carefully observing the motions of these stars, astronomers can glean important information about their masses.

### Binary Stars

A pair of stars located at nearly the same position in the night sky is called a **double star.** The Anglo-German astronomer William Herschel made the first organized search for such pairs. Between 1782 and 1821, he published three catalogs listing more than 800 double stars. Late in the nineteenth century, his son, John Herschel, discovered 10,000 more doubles. Some of these double stars are **optical double stars,** which are two stars that lie along nearly the same line of sight but are actually at very different distances from us. But many double stars are true **binary stars,** or **binaries**—pairs of stars that actually orbit each other. Figure 17-19 shows an example of this orbital motion.



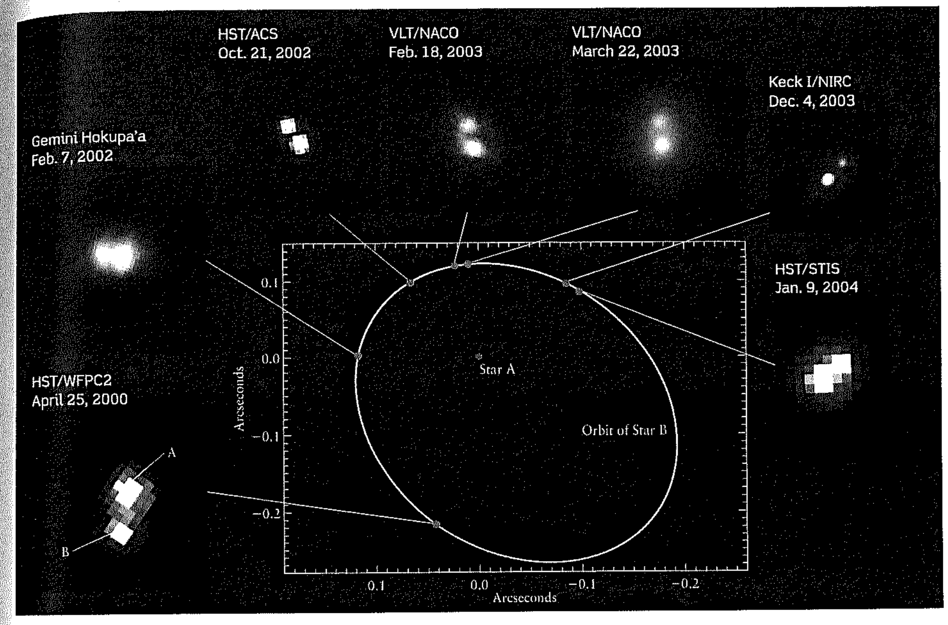



**Figure 17-19** R I V U X G

A Binary Star System As seen from Earth, the two stars that make up the binary system called 2MASSW J0746425+2000321 are separated by less than 1/3 arcsecond. The images surrounding the center diagram show the relative positions of the two stars over a four-year period. These images were made by the Hubble Space Telescope (HST), the European Southern Observatory's Very Large Telescope (VLT), and Keck I and Gemini North in Hawaii (see Figure 6-16). For simplicity, the diagram shows one star as remaining stationary; in reality, both stars move around their common center of mass. (H. Bouy et al., MPE and ESO)

When astronomers can actually see the two stars orbiting each other, a binary is called a **visual binary.** By observing the binary over an extended period, astronomers can plot the orbit that one star appears to describe around the other, as shown in the center diagram in Figure 17-19.

In fact, *both* stars in a binary system are in motion. They orbit each other because of their mutual gravitational attraction, and their orbital motions obey Kepler's third law as formulated by Isaac Newton (see Section 4-7 and Box 4-4). This law can be written as follows:

Kepler's third law for binary star systems

$$M_1 + M_2 = \frac{a^3}{P^2}$$

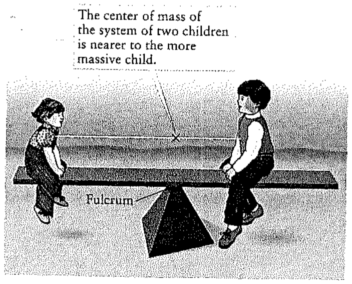
$M_1$, $M_2$ = masses of two stars in binary system, in solar masses

$a$ = semimajor axis of one star's orbit around the other, in AU

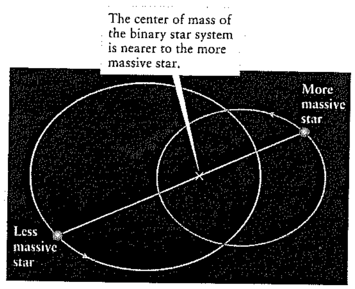
$P$ = orbital period, in years

Here $a$ is the semimajor axis of the elliptical orbit that one star appears to describe around the other, plotted as in the center diagram in Figure 17-19. As this equation indicates, if we can measure this semimajor axis ($a$) and the orbital period ($P$), we can learn something about the masses of the two stars.

In principle, the orbital period of a visual binary is easy to determine. All you have to do is see how long it takes for the two stars to revolve once about each other. The two stars shown in Figure 17-19 are relatively close, about 2.5 AU on average, and their orbital period is only 10 years. Many binary systems have




(a) A "binary system" of two children




(b) A binary star system

**Figure 17-20**

Center of Mass in a Binary Star System (a) A seesaw balances if the fulcrum is at the center of mass of the two children. (b) The members of a binary star system orbit around the center of mass of the two stars. Although their elliptical orbits cross each other, the two stars are always on opposite sides of the center of mass and thus never collide.

much larger separations, however, and the period may be so long that more than one astronomer's lifetime is needed to complete the observations.

Determining the semimajor axis of an orbit can also be a challenge. The *angular* separation between the stars can be determined by observation. To convert this angle into a physical distance between the stars, we need to know the distance between the binary and Earth. This distance can be found from parallax measurements or by using spectroscopic parallax. The astronomer must also take into account how the orbit is tilted to our line of sight.

Once both $P$ and $a$ have been determined, Kepler's third law can be used to calculate $M_1 + M_2$, the sum of the masses of the two stars in the binary system. But this analysis tells us nothing about the *individual* masses of the two stars. To obtain these masses, more information about the motions of the two stars is needed.

Each of the two stars in a binary system actually moves in an elliptical orbit about the **center of mass** of the system. Imagine two children sitting on opposite ends of a seesaw (Figure 17-20*a*). For the seesaw to balance properly, they must position themselves so that their center of mass—an imaginary point that lies along a line connecting their two bodies—is at the fulcrum, or pivot point of the seesaw. If the two children have the same mass, the center of mass lies midway between them, and they should sit equal distances from the fulcrum. If their masses are different, the center of mass is closer to the heavier child.

Just as the seesaw naturally balances at its center of mass, the two stars that make up a binary system naturally orbit around their center of mass (Figure 17-20*b*). The center of mass always lies along the line connecting the two stars and is closer to the more massive star.

The center of mass of a visual binary is located by plotting the separate orbits of the two stars, as in Figure 17-20*b*, using the background stars as reference points. The center of mass lies at the common focus of the two elliptical orbits. Comparing the relative sizes of the two orbits around the center of mass yields the ratio of the two stars' masses, $M_1/M_2$. The sum $M_1 + M_2$ is already known from Kepler's third law, so the individual masses of the two stars can then be determined.

## Main-Sequence Masses and the Mass-Luminosity Relation

Years of careful, patient observations of binaries have slowly yielded the masses of many stars. As the data accumulated, an important trend began to emerge: For main-sequence stars, there is a direct correlation between mass and luminosity. The more massive a main-sequence star, the more luminous it is. Figure 17-21 depicts this **mass-luminosity relation** as a graph. The range of stellar masses extends from less than 0.1 of a solar mass to more than 50 solar masses. The Sun's mass lies between these extremes.

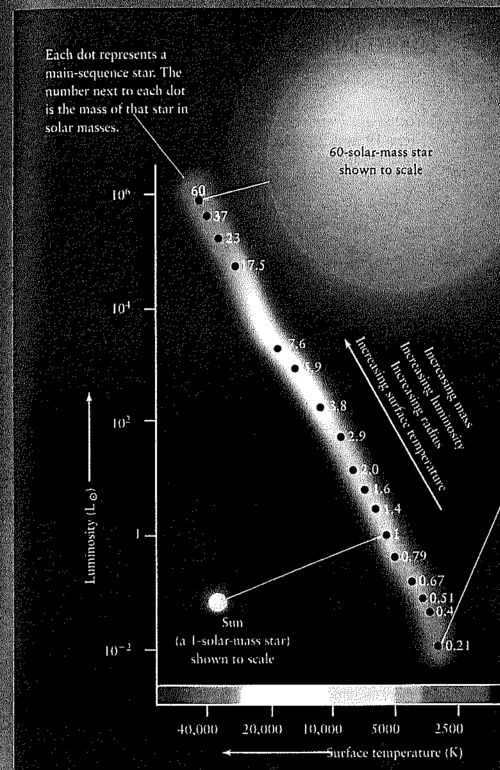
The *Cosmic Connections* figure on the next page depicts the mass-luminosity relation for main-sequence stars on an H-R diagram. This figure shows the main sequence on an H-R diagram is a progression in mass as well as in luminosity and surface temperature. The hot, bright, bluish stars in the upper left corner of an H-R diagram are the most massive main-sequence stars. Likewise, the dim, cool, reddish stars in the lower right corner of an H-R diagram are the least massive. Main-sequence stars of intermediate temperature and luminosity also have intermediate masses.

The mass of a main-sequence star also helps determine its radius. Referring back to Figure 17-15*b*, we see that if we move along the main sequence from low luminosity to high luminosity, the radius of the star increases. Thus, we have the following general rule for main-sequence stars:

***The greater the mass of a main-sequence star, the greater its luminosity, its surface temperature, and its radius.***

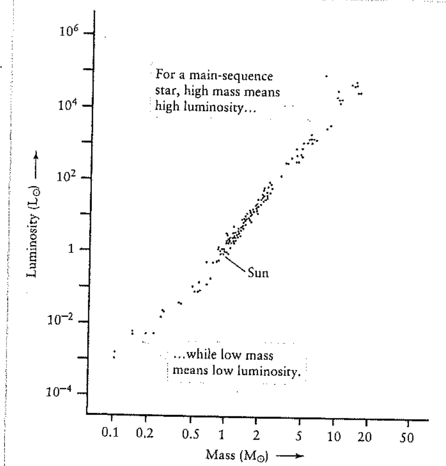
# COSMIC CONNECTIONS The Main Sequence and Masses

The main sequence is an arrangement of stars according to their mass. The most massive main-sequence stars have the greatest luminosity, greatest radius, and greatest surface temperature. These characteristics are consequences of the behavior of thermonuclear reactions at the core of a main-sequence star.




- A star with 60 solar masses has much higher pressure and temperature at its core than does the Sun.
- This heat and pressure cause thermonuclear reactions in the core to occur much more rapidly and release energy at a much faster rate — 790,000 times faster than in the Sun.
- Energy is emitted from the star's surface at the same rate that it is released in the core, so the star has 790,000 times the Sun's luminosity.
- The tremendous rate of energy release also heats the star's interior tremendously, increasing the star's internal pressure. This inflates the star to 15 times the Sun's radius.
- The star's surface must be at a high temperature (about 44,500 K) in order for it to emit energy into space at such a rapid rate.

- A star with 0.21 solar mass has much lower pressure and temperature at its core than does the Sun.
- This causes thermonuclear reactions in the core to occur much more slowly and release energy at a much slower rate — 0.011 times as fast as in the Sun.
- Energy is emitted from the star's surface at the same rate that it is released in the core, so the star has 0.011 of the Sun's luminosity.
- The low rate of energy release supplies relatively little heat to the star's interior, so the star's internal pressure is low. Hence the star's radius is only 0.33 times the Sun's radius.
- The star's surface need be at only a low temperature (about 3200 K) to emit energy into space at such a relatively slow rate.




## Figure 17-21

The Mass-Luminosity Relation For main-sequence stars, there is a direct correlation between mass and luminosity—the more massive a star, the more luminous it is. A main-sequence star of mass 10 $M_\odot$ (that is, 10 times the Sun's mass) has roughly 3000 times the Sun's luminosity (3000 $L_\odot$); one with 0.1 $M_\odot$ has a luminosity of only about 0.001 $L_\odot$.

### Mass and Main-Sequence Stars

Why is mass the controlling factor in determining the properties of a main-sequence star? The answer is that all main-sequence stars are objects like the Sun, with essentially the same chemical composition as the Sun but with different masses. Like the Sun, all main-sequence stars shine because thermonuclear reactions at their cores convert hydrogen to helium and release energy. The greater the total mass of the star, the greater the pressure and temperature at the core, the more rapidly thermonuclear reactions take place in the core, and the greater the energy output—that is, the luminosity—of the star. In other words, the greater the mass of a main-sequence star, the greater its luminosity. This statement is just the mass-luminosity relation, which we can now recognize as a natural consequence of the nature of main-sequence stars.

Like the Sun, main-sequence stars are in a state of both hydrostatic equilibrium and thermal equilibrium. Calculations using models of a main-sequence star's interior (like the solar models we discussed in Section 16-2) show that to maintain equilibrium, a more massive star must have a larger radius and a higher surface temperature. This result is just what we see when we plot the curve of the main sequence on an H-R diagram (see Figure 17-15*b*). As you move up the main sequence from less massive stars (at the lower right in the H-R diagram) to more massive stars (at the upper left), the radius and surface temperature both increase.

Calculations using hydrostatic and thermal equilibrium also show that if a star's mass is less than about 0.08$M_\odot$, the core pressure and temperature are too low for thermonuclear reactions to take place. The "star" is then a brown dwarf. Brown dwarfs also follow a mass-luminosity relation: The greater the mass, the faster the brown dwarf contracts because of its own gravity, the more rapidly it radiates energy into space, and, hence, the more luminous the brown dwarf is.

**CAUTION!** The mass-luminosity relation we have discussed applies to main-sequence stars only. There are *no* simple mass-luminosity relations for giant, supergiant, or white dwarf stars. Why these stars lie where they do on an H-R diagram will become apparent when we study the evolution of stars in Chapters 19 and 20. We will find that main-sequence stars evolve into giant and supergiant stars, and that some of these eventually end their lives as white dwarfs.

## 17-10 Spectroscopy makes it possible to study binary systems in which the two stars are close together

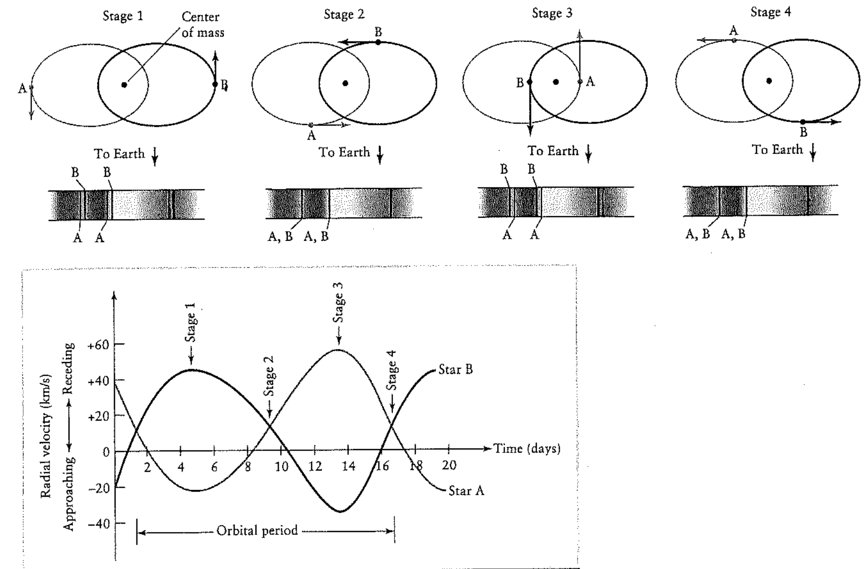
We have described how the masses of stars can be determined from observations of visual binaries, in which the two stars can be distinguished from each other. But if the two stars in a binary system are too close together, the images of the two stars can blend to produce the semblance of a single star. Happily, in many cases we can use spectroscopy to decide whether a seemingly single star is in fact a binary system. Spectroscopic observations of binaries provide additional useful information about star masses.

Some binaries are discovered when the spectrum of a star shows incongruous spectral lines. For example, the spectrum of what appears to be a single star may include both strong hydrogen lines (characteristic of a type A star) and strong absorption bands of titanium oxide (typical of a type M star). Because a single star cannot have the differing physical properties of these two spectral types, such a star must actually be a binary system that is too far away for us to resolve its individual stars. A binary system detected in this way is called a **spectrum binary**.

Other binary systems can be detected using the Doppler effect. If a star is moving toward Earth, its spectral lines are displaced toward the short-wavelength (blue) end of the spectrum. Conversely, the spectral lines of a star moving away from us are shifted toward the long-wavelength (red) end of the spectrum. The upper portion of Figure 17-22 applies these ideas to a hypothetical binary star system with an orbital plane that is edge-on to our line of sight.

As the two stars move around their orbits, they periodically approach and recede from us. Hence, the spectral lines of the two stars are alternately blueshifted and redshifted. The two stars in this hypothetical system are so close together that they appear through a telescope as a single star with a single spectrum.






## Figure 17-22

Radial Velocity Curves The lower graph displays the radial velocity curves of the binary system HD 171978. The drawings at the top indicate the positions of the stars (labeled A and B) and the spectra of the binary at four selected moments (stages 1, 2, 3, and 4) during an orbital period. Note that at stages 1 and 3, the Doppler effect splits apart the absorption lines from stars A and B; compare with Figure 17-23. The entire binary star system is moving away from us at 12 km/s, which is why the entire pattern of radial velocity curves is displaced upward from the zero-velocity line.

Because one star shows a blueshift while the other is showing a redshift, the spectral lines of the binary system appear to split apart and rejoin periodically. Stars whose binary character is revealed by such shifting spectral lines are called **spectroscopic binaries.**

### Exploring Spectroscopic Binary Stars

**Stars in close binary systems move so rapidly that we can detect their motion using the Doppler effect**

To analyze a spectroscopic binary, astronomers measure the wavelength shift of each star's spectral lines and use the Doppler shift formula (introduced in Section 5-9 and Box 5-6) to determine the **radial velocity** of each star—that is, how fast and in what direction it is moving along our line of sight. The lower portion of Figure 17-22 shows a graph of the radial velocity versus time, called a **radial velocity curve,** for the binary system HD 171978. Each of the two stars alternately approaches and recedes as it orbits around the center of mass. The pattern of the curves repeats every 15 days, which is the orbital period of the binary.

Figure 17-23 shows two spectra of the spectroscopic binary κ (kappa) Arietis taken a few days apart. In Figure 17-23*a*, two sets of spectral lines are visible, offset slightly in opposite directions from the normal positions of these lines. This corresponds to stage 1 or stage 3 in Figure 17-22; one of the orbiting stars is moving toward Earth and has its spectral lines blueshifted, and the other star is moving away from Earth and has its lines redshifted. A few days later, the stars have progressed along their orbits so that neither star is moving toward or away from Earth, corresponding to stage 2 or stage 4 in Figure 17-22. At this time there are no Doppler shifts, and the spectral lines of both stars are at the same positions. That is why only one set of spectral lines appears in Figure 17-23*b*.

It is important to emphasize that the Doppler effect applies only to motion along the line of sight. Motion perpendicular to

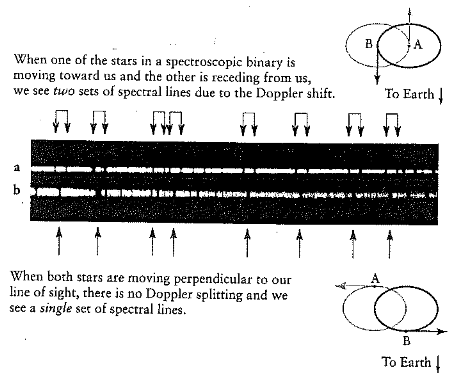



**Figure 17-23** R I V U X G

A Spectroscopic Binary The visible-light spectrum of the double-line spectroscopic binary κ (kappa) Arietis has spectral lines that shift back and forth as the two stars revolve about each other. (Lick Observatory)

the line of sight does not affect the observed wavelengths of spectral lines. Hence, the ideal orientation for a spectroscopic binary is to have the stars orbit in a plane that is edge-on to our line of sight. (By contrast, a *visual* binary is best observed if the orbital plane is face-on to our line of sight.) For the Doppler shifts to be noticeable, the orbital speeds of the two stars should be at least a few kilometers per second.

The binaries depicted in Figures 17-22 and 17-23 are called *double-line* spectroscopic binaries, because the spectral lines of both stars in the binary system can be seen. Most spectroscopic binaries, however, are *single-line* spectroscopic binaries, in which one of the stars is so dim that its spectral lines cannot be detected. The star is obviously a binary, however, because its spectral lines shift back and forth, thereby revealing the orbital motions of two stars about their center of mass.

As for visual binaries, spectroscopic binaries allow astronomers to learn about stellar masses. From a radial velocity curve, one can find the *ratio* of the masses of the two stars in a binary. The *sum* of the masses is related to the orbital speeds of the two stars by Kepler's laws and Newtonian mechanics. If both the ratio of the masses and their sum are known, the individual masses can be determined using algebra. However, determining the sum of the masses requires that we know how the binary orbits are tilted from our line of sight. This is because the Doppler shifts reveal only the radial velocities of the stars rather than their true orbital speeds. This tilt is often impossible to determine, because we cannot see the individual stars in the binary. Thus, the masses of stars in spectroscopic binaries tend to be uncertain.

There is one important case in which we can determine the orbital tilt of a spectroscopic binary. If the two stars are observed to eclipse each other periodically, then we must be viewing the orbit nearly edge-on. As we will see next, individual stellar masses—as well as other useful data—can be determined if a spectroscopic binary also happens to be such an *eclipsing* binary.

## 17-11 Light curves of eclipsing binaries provide detailed information about the two stars

Some binary systems are oriented so that the two stars periodically eclipse each other as seen from Earth. These **eclipsing binaries** can be detected even when the two stars cannot be resolved visually as two distinct images in the telescope. The apparent brightness of the image of the binary dims briefly each time one star blocks the light from the other.

Using a sensitive detector at the focus of a telescope, an astronomer can measure the incoming light intensity quite accurately and create a **light curve** (Figure 17-24). The shape of the light curve for an eclipsing binary reveals at a glance whether the eclipse is partial or total (compare Figures 17-24*a* and 17-24*b*). Figure 17-24*d* shows an observation of a binary system undergoing a total eclipse.

**Eclipsing binaries can reveal the sizes and shapes of stars**

In fact, the light curve of an eclipsing binary can yield a surprising amount of information. For example, the ratio of the surface temperatures can be determined from how much their combined light is diminished when the stars eclipse each other. Also, the duration of a mutual eclipse tells astronomers about the relative sizes of the stars and their orbits.

If the eclipsing binary is also a double-line spectroscopic binary, an astronomer can calculate the mass and radius of each star from the light curves and the velocity curves. Unfortunately, very few binary stars are of this ideal type. Stellar radii determined in this way agree well with the values found using the Stefan-Boltzmann law, as described in Section 17-6.

The shape of a light curve can reveal many additional details about a binary system. In some binaries, for example, the gravitational pull of one star distorts the other, much as the Moon distorts Earth's oceans in producing tides (see Figure 4-26). Figure 17-24*c* shows how such tidal distortion gives the light curve a different shape than in Figure 17-24*b*.

Information about stellar atmospheres can also be derived from light curves. Suppose that one star of a binary is a luminous main-sequence star and the other is a bloated red giant. By observing exactly how the light from the bright main-sequence star is gradually cut off as it moves behind the edge of the red giant during the beginning of an eclipse, astronomers can infer the pressure and density in the upper atmosphere of the red giant.

Binary systems are tremendously important because they enable astronomers to measure stellar masses as well as other key properties of stars. In the next several chapters, we will use this information to help us piece together the story of *stellar evolution*—how stars are born, evolve, and eventually die.



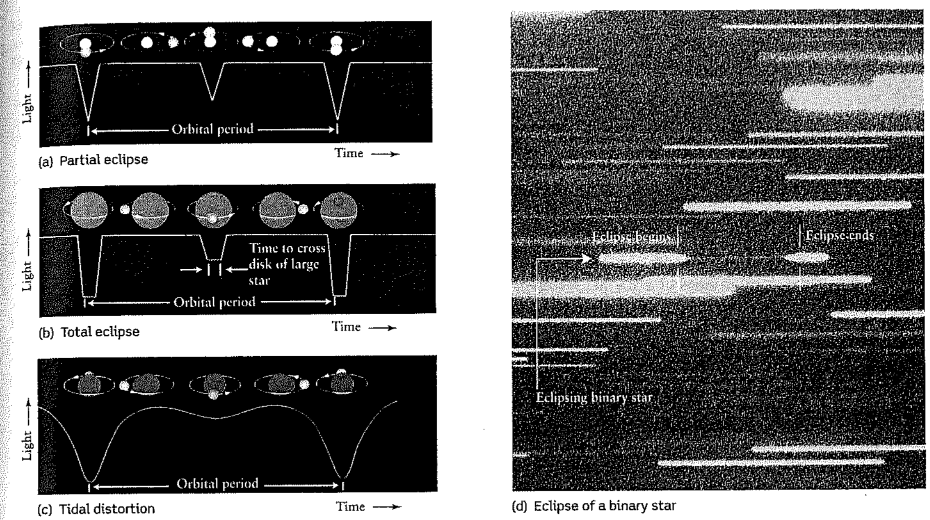



**Figure 17-24** R I V U X G

Representative Light Curves of Eclipsing Binaries **(a)**, **(b)**, **(c)** The shape of the light curve of an eclipsing binary can reveal many details about the two stars that make up the binary. **(d)** This image shows the binary star NN Serpens (indicated by the arrow) undergoing a total eclipse. The telescope was moved during the exposure so that the sky drifted slowly from left to right across the field of view. During the 10.5-minute duration of the eclipse, the dimmer star of the binary system (an M6 main-sequence star) passed in front of the other, more luminous star (a white dwarf). The binary became so dim that it almost disappeared. (European Southern Observatory)

## Key Words

*Terms preceded by an asterisk (*) are discussed in the Boxes.*

absolute magnitude, p. 441
apparent brightness (brightness), p. 437
apparent magnitude, p. 440
binary star (binary), p. 456
brown dwarf, p. 449
center of mass, p. 458
color ratio, p. 445
*distance modulus, p. 444
double star, p. 456
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## Key Ideas

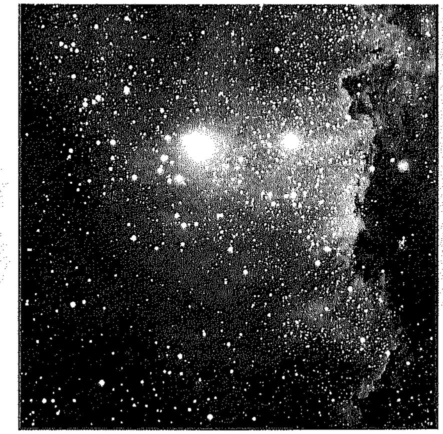
**Measuring Distances to Nearby Stars:** Distances to the nearer stars can be determined by parallax, the apparent shift of a star against the background stars observed as Earth moves along its orbit.

• Parallax measurements made from orbit, above the blurring effects of the atmosphere, are much more accurate than those made with Earth-based telescopes.

• Stellar parallaxes can only be measured for stars within a few hundred parsecs.

**The Inverse-Square Law:** A star's luminosity (total light output), apparent brightness, and distance from Earth are related by the

# 18
# The Birth of Stars



R I V U X G
A region of star formation about 1400 pc (4000 ly) from Earth in the southern constellation Ara (the Altar). (European Southern Observatory)

The stars that illuminate our nights seem eternal and unchanging. But this permanence is an illusion. Each of the stars visible to the naked eye shines due to thermonuclear reactions and has only a finite amount of fuel available for these reactions. Hence, stars cannot last forever: They form from material in interstellar space, evolve over millions or billions of years, and eventually die. In this chapter our concern is with how stars are born and become part of the main sequence.

Stars form within cold, dark clouds of gas and dust that are scattered abundantly throughout our Galaxy. One such cloud appears as a dark area on the far right-hand side of the photograph at the top of this page. Perhaps a dark cloud like this encounters one of the Galaxy's spiral arms, or perhaps a supernova detonates nearby. From the shock of events like these, the cloud begins to contract under the pull of gravity, forming protostars—the fragments that will one day become stars. As a protostar develops, its internal pressure builds and its temperature rises. In time, hydrogen fusion begins, and a star is born. The hottest, bluest, and brightest young stars, like those in the accompanying image, emit ultraviolet radiation that excites the surrounding interstellar gas. The result is a beautiful glowing nebula, which typically has the red color characteristic of excited hydrogen (as shown in the photograph).

In Chapters 19 and 20 we will see how stars mature and grow old. Some even blow themselves apart in death throes that enrich interstellar space with the material for future generations of stars. Thus, like the mythical phoenix, new stars arise from the ashes of the old.

## 18-1 Understanding how stars evolve requires observation as well as ideas from physics

Over the past several decades, astronomers have labored to develop an understanding of **stellar evolution,** that is, how stars are born, live their lives, and finally die. Our own Sun provides evidence that stars are not permanent. The energy radiated by the Sun comes from thermonuclear reactions in its core, which consume $6 \times 10^{11}$ kg of hydrogen each second and convert it into helium (see Section 16-1). While the amount of hydrogen in the Sun's core is vast, it is not infinite; therefore, the Sun cannot always have been shining, nor can it continue to shine forever. The same is true for all other main-sequence stars, which are fundamentally the same kinds of objects as the Sun but with different masses (see Section 17-9). Thus, stars must have a beginning as well as an end.

**Stars consume the material of which they are made, and so cannot last forever**

## Learning Goals

*By reading the sections of this chapter, you will learn*

18-1 How astronomers have pieced together the story of stellar evolution

18-2 What interstellar nebulae are and what they are made of

18-3 What happens as a star begins to form

18-4 The stages of growth from young protostars to main-sequence stars

18-5 How stars gain and lose mass during their growth

18-6 What insights star clusters add to our understanding of stellar evolution

18-7 Where new stars form within galaxies

18-8 How the death of old stars can trigger the birth of new stars

Stars last very much longer than the lifetime of any astronomer—indeed, far longer than the entire history of human civilization. Thus, it is impossible to watch a single star go through its formation, evolution, and eventual demise. Rather, astronomers have to piece together the evolutionary history of stars by studying different stars at different stages in their life cycles.

ANALOGY To see the magnitude of this task, imagine that you are a biologist from another planet who sets out to understand the life cycles of human beings. You send a spacecraft to fly above Earth and photograph humans in action. Unfortunately, the spacecraft fails after collecting only 20 seconds of data, but during that time its sophisticated equipment sends back observations of thousands of different humans. From this brief snapshot of life on Earth—only $10^{-8}$ (a hundred-millionth) of a typical human lifetime—how would you decide which were the young humans and which were the older ones? Without a look inside our bodies to see the biological processes that shape our lives, could you tell how humans are born and how they die? And how could you deduce the various biological changes that humans undergo as they age?

Astronomers, too, have data spanning only a tiny fraction of any star's lifetime. A star like the Sun can last for about $10^{10}$ years, whereas astronomers have been observing stars in detail for only about a century—as in our analogy, roughly $10^{-8}$ of the life span of a typical star. Astronomers are also frustrated by being unable to see the interiors of stars. For example, we cannot see the thermonuclear reactions that convert hydrogen into helium. But astronomers have an advantage over the biologist in our story. Unlike humans, stars are made of relatively simple substances, primarily hydrogen and helium, that are found almost exclusively in the form of gases. Of the three phases of matter—gas, liquid, and solid—gases are by far the simplest to understand.

Astronomers use our understanding of gases to build theoretical models of the interiors of stars, like the model of the Sun we saw in Section 16-2. Models help to complete the story of stellar evolution. In fact, like all great dramas, the story of stellar evolution can be regarded as a struggle between two opposing and unyielding forces: Gravity continually tries to make a star shrink, while the star's internal pressure tends to make the star expand. When these two opposing forces are in balance, the star is in a state of hydrostatic equilibrium (see Figure 16-2).

But what happens when changes within the star cause either pressure or gravity to predominate? The star must then either expand or contract until it reaches a new equilibrium. In the process, it will change not only in size but also in luminosity and color.

In the following chapters, we will find that giant and supergiant stars are the result of pressure gaining the upper hand over gravity. Both giants and supergiants turn out to be aging stars that have become tremendously luminous and ballooned to hundreds or thousands of times their previous size. White dwarfs, by contrast, are the result of the balance tipping in gravity's favor. These dwarfs are even older stars that have collapsed to a fraction of the size they had while on the main sequence. In this chapter, however, we will see how the opposing influences of gravity and pressure explain the birth of stars. We start our journey within the diffuse clouds of gas and dust that permeate our galaxy.

## 18-2 Interstellar gas and dust pervade the galaxy

**Different types of nebulae emit, absorb, or reflect light**

Where do stars come from? As we saw in Section 8-4, our Sun condensed from a solar nebula, a collection of gas and dust in interstellar space. Observations suggest that other stars originate in a similar way (see Figure 8-8). To understand the formation of stars, we must first understand the nature of the interstellar matter from which the stars form.

### Nebulae and the Interstellar Medium

At first glance, the space between the stars seems to be empty. On closer inspection, we find that it is filled with a thin gas laced with microscopic dust particles. This combination of gas and dust is called the **interstellar medium.** Evidence we'll discuss for this medium includes interstellar clouds of various types, curious lines in the spectra of binary star systems, and an apparent dimming and reddening of distant stars.

You can see evidence for the interstellar medium with the naked eye. Look carefully at the constellation Orion (Figure 18-1*a*), visible on winter nights in the northern hemisphere and summer nights in the southern hemisphere. While most of the stars in the constellation appear as sharply defined points of light, the middle "star" in Orion's sword has a fuzzy appearance. This fuzziness becomes more obvious with binoculars or a telescope. As Figure 18-1*b* shows, this "star" is actually not a star at all, but the Orion Nebula—a cloud in interstellar space. Any interstellar cloud is called a **nebula** (plural **nebulae**) or **nebulosity.**

### Emission Nebulae: Clouds of Excited Gas

The Orion Nebula emits its own light, with the characteristic emission line spectrum of a hot, thin gas. For this reason it is called an **emission nebula.** Many emission nebulae can be seen with a small telescope. Figure 18-2 shows some of these nebulae in a different part of the constellation Orion. Emission nebulae are direct evidence of gas atoms in the interstellar medium.

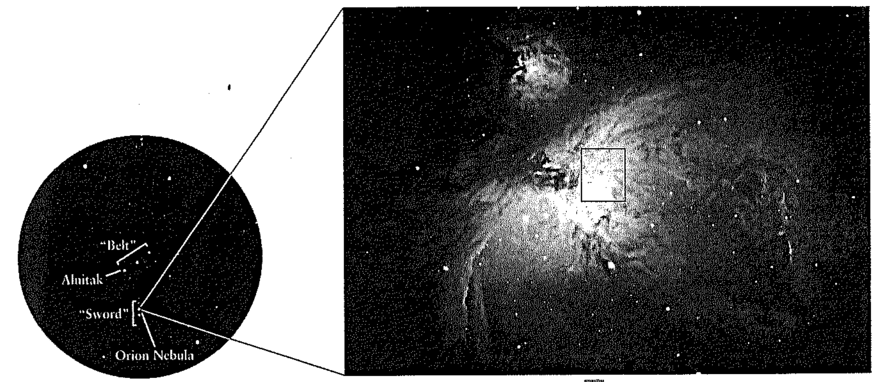
Typical emission nebulae have masses that range from about 100 to about 10,000 solar masses. Because this mass is spread over a huge volume that is light-years across, the density is quite low by Earth standards, only a few thousand hydrogen atoms per cubic centimeter. (By comparison, the air you are breathing contains more than $10^{19}$ atoms per $cm^3$.)

Emission nebulae are found near hot, luminous stars of spectral types O and B. Such stars emit copious amounts of ultraviolet radiation. When atoms in the nearby interstellar gas absorb these energetic ultraviolet photons, the atoms become ionized. Indeed, emission nebulae are composed primarily of ionized hydrogen atoms, that is, free protons (hydrogen nuclei) and electrons. Astronomers use the notation H I for neutral, un-ionized hydrogen atoms and H II for ionized hydrogen atoms, which is why emission nebulae are also called **H II regions.**

H II regions emit visible light when some of the free protons and electrons get back together to form hydrogen atoms, a process called **recombination** (Figure 18-3). When an atom forms by recombination, the electron is typically captured into a high-energy orbit. As the electron cascades downward through the atom's energy levels toward the ground state, the atom emits photons with lower energies and longer wavelengths than the photons that originally caused the ionization. Particularly important is the transition from $n = 3$ to $n = 2$. It produces $H_\alpha$ photons with a wavelength of 656 nm, in the red portion of the visible spectrum (see Section 5-8, especially Figure 5-23*b*). These photons give H II regions their distinctive reddish color.

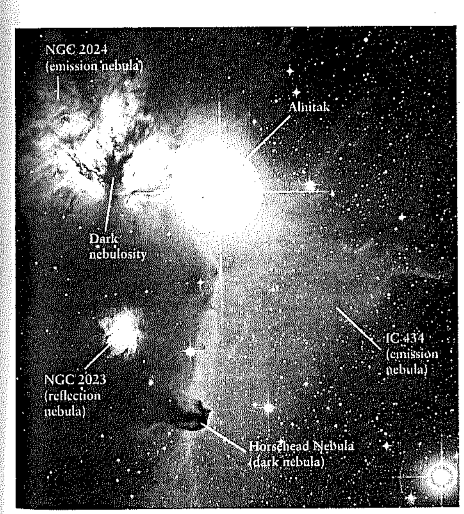
For each high-energy, ultraviolet photon absorbed by a hydrogen atom to ionize it, several photons of lower energy are emitted when a proton and electron recombine. As Box 18-1 describes, a similar effect takes place in a fluorescent lightbulb. In this sense, H II regions are immense fluorescent light fixtures!



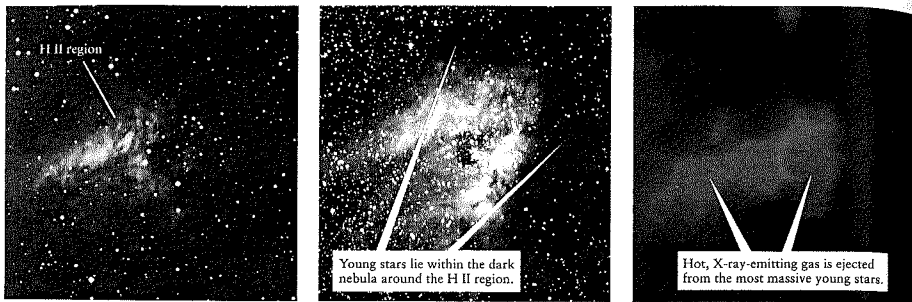



(a) A wide-angle view of Orion (b) A closeup of the Orion Nebula R I V U X G

WEB LINK 18.2 **Figure 18-1**
The Orion Nebula (a) The middle "star" of the three that make up Orion's sword is actually an interstellar cloud called the Orion Nebula. (b) The nebula is about 450 pc (1500 ly) from Earth and contains about 300 solar masses of material. Within the area shown by the box are four hot, massive stars called the Trapezium. They produce the ultraviolet light that makes the nebula glow. (a: R. C. Mitchell, Central Washington University; b: Anglo-Australian Observatory)




WEB LINK 18.3 **Figure 18-2** R I V U X G
Emission, Reflection, and Dark Nebulae in Orion A variety of different nebulae appear in the sky around Alnitak, the easternmost star in Orion's belt (see Figure 18-1a). All the nebulae lie approximately 500 pc (1600 ly) from Earth. They are actually nowhere near Alnitak, which is only 250 pc (820 ly) distant. This photograph shows an area of the sky about 1.5° across. (Royal Observatory, Edinburgh)




(a) Visible-light image R I V U X G (b) False-color infrared image R I V U X G (c) False-color X-ray image R I V U X G

**Figure 18-13**
Mass Loss from Young, Massive Stars **(a)** The Omega Nebula, also known as M17, is a region of star formation in the constellation Sagittarius about 1700 pc (5500 ly) from Earth. **(b)** This infrared image allows us to see through dust, revealing recently formed stars that cannot be seen in (a). **(c)** The most massive young stars eject copious amounts of hot gas. Red indicates X-ray emission from gas at a temperature of $1.5 \times 10^6$ K; blue indicates even hotter gas at a temperature of $7 \times 10^6$ K. Astronomers do not see such X-ray emission from the Orion Nebula (Figure 18-1), which has many young stars but very few massive ones. (a: Palomar Observatory DSS; b: 2MASS/UMass/IPAC-Caltech/NASA/NSF; c: NASA/CXC/PSU/L. Townsley et al.)

mass. This process is called **accretion,** and the disk of material being added to the protostar in this way is called a **circumstellar accretion disk.** Figure 18-15 is an edge-on view of a circumstellar accretion disk, showing two oppositely directed jets emanating from a point at or near the center of the disk (where the protostar is located).

What causes some of the material in the disk to be blasted outward in a pair of jets? One model involves the magnetic field of the dark nebula in which the star forms (Figure 18-16). As material in the circumstellar accretion disk falls inward, it drags the magnetic field lines along with it. (We saw in Section 16-9 how a similar mechanism may explain the Sun's 22-year cycle.) Parts of the disk at different distances from the central protostar orbit at different speeds, and this can twist the magnetic field lines into two helix shapes, one on each side of the disk. The helices then act as channels that guide infalling material away from the protostar, forming two opposing jets.

Many astronomers suspect that interactions among the protostar, the accretion disk, and the jets help to slow the protostar's rotation. If so, this would explain why main-sequence stars generally spin much more slowly than protostars of the same final mass.

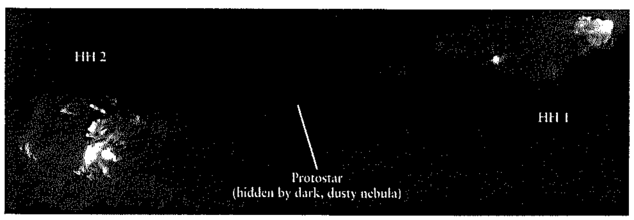



**Figure 18-14** R I V U X G
Bipolar Outflow and Herbig-Haro Objects The two bright knots of glowing, ionized gas called HH 1 and HH 2 are Herbig-Haro objects. They are created when fast-moving gas ejected from a protostar slams into the surrounding interstellar medium, heating the gas to high temperature. HH 1 and HH 2 are 0.34 parsec (1.1 light-year) apart and lie 470 pc (1500 ly) from Earth in the constellation Orion. (J. Hester, the WFPC-2 Investigation Definition Team, and NASA)



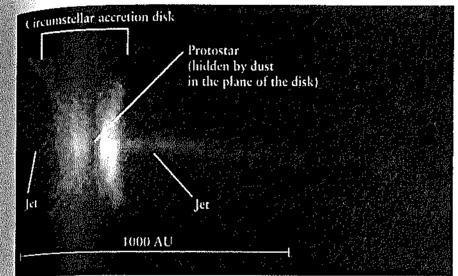



**Figure 18-15** R I V U X G
A Circumstellar Accretion Disk and Jets This false-color image shows a star surrounded by an accretion disk, which we see nearly edge-on. Red denotes emission from ionized gas, while green denotes starlight scattered from dust particles in the disk. The midplane of the accretion disk is so dusty and opaque that it appears dark. Two oppositely directed jets flow away from the star, perpendicular to the disk and along the disk's rotation axis. This star lies 140 pc (460 ly) from Earth. (C. Burrows, the WFPC-2 Investigation Definition Team, and NASA)

In the 1990s, astronomers using the Hubble Space Telescope discovered many examples of disks around newly formed stars in the Orion Nebula (see Figure 18-1), one of the most prominent star-forming regions in the northern sky. Figure 8-8 shows a number of these **protoplanetary disks,** or **proplyds,** that surround young stars within the nebula. As the name suggests, protoplanetary disks are thought to contain the material from which planets form around stars. They are what remains of a circumstellar accretion disk after much of the material has either fallen onto the star or been ejected by bipolar outflows.

Not all stars are thought to form protoplanetary disks; the exceptions probably include stars with masses in excess of about 3 $M_{\odot}$, as well as many stars in binary systems. But surveys of the Orion Nebula show that disks are found around most young, low-mass stars. Thus, disk formation may be a natural stage in the birth of many stars.

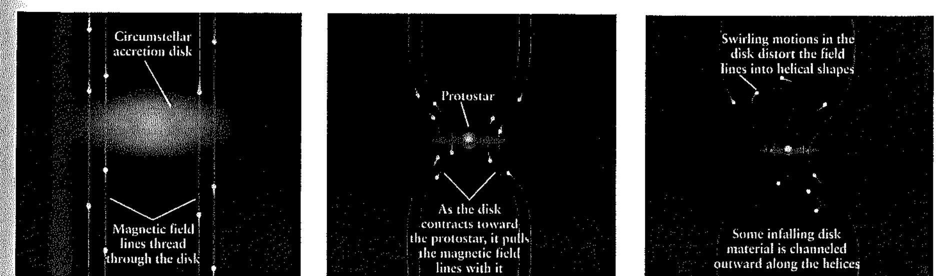
## 18-6 Young star clusters give insight into star formation and evolution

Dark nebulae contain tens or hundreds of solar masses of gas and dust, enough to form many stars. As a consequence, these nebulae tend to form groups or clusters of young stars. One such cluster is M16, shown in Figure 18-17; another is NGC 6520, visible in Figure 18-4.

### Star Clusters as Evolutionary Laboratories

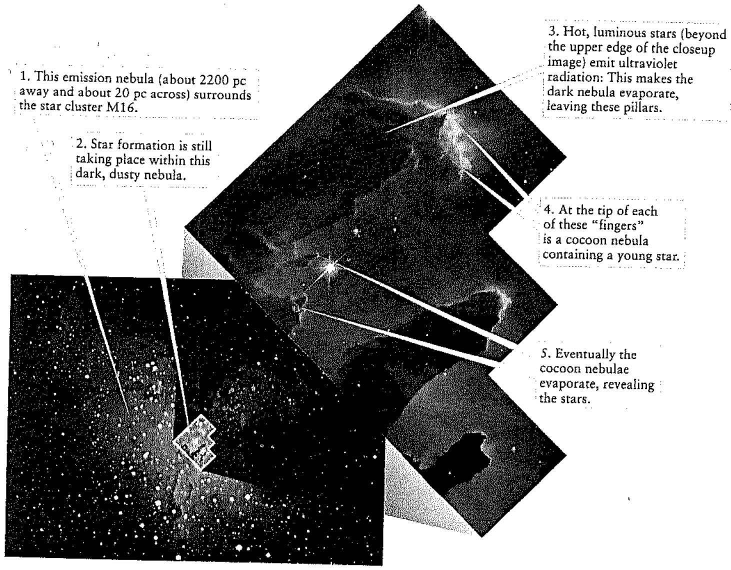
In addition to being objects of great natural beauty, star clusters give us a unique way to compare the evolution of different stars. That's because clusters typically include stars with a range of different masses, all of which began to form out of the parent nebula at roughly the same time.

ANALOGY A foot race is a useful way to compare the performance of sprinters because all the competitors start the race simultaneously. A young star cluster gives us the same kind of opportunity to compare the evolution of stars of different masses that all began to form roughly simultaneously. Unlike a foot




(a) (b) (c)

**Figure 18-16**
A Magnetic Model for Bipolar Outflow **(a)** Observations suggest that circumstellar accretion disks are threaded by magnetic field lines, as shown here. **(b), (c)** The contraction and rotation of the disk make the magnetic field lines distort and twist into helices. These helices steer some of the disk material into jets that stream perpendicular to the plane of the disk, as in Figure 18-15. (Adapted from Alfred T. Kamajian/Thomas P. Ray, "Fountain of Youth: Early Days in the Life of a Star," *Scientific American*, August 2000)




WEB LINK 18.1 ANIMATION 18.3

**Figure 18-17** R I **V** U X G

A Star Cluster with an H II Region The star cluster M16 is thought to be no more than 800,000 years old, and star formation is still taking place within adjacent dark, dusty globules. The inset shows three dense, cold pillars of gas and dust silhouetted against the glowing background of the red emission nebula (called the Eagle Nebula for its shape). The pillar at the upper left extends about 0.3 parsec (1 light-year) from base to tip, and each of its "fingers" is somewhat broader than our entire solar system. (Anglo-Australian Observatory; J. Hester and P. Scowen, Arizona State University; NASA)

race, however, the entire "race" of stellar evolution in a single cluster happens too slowly for us to observe; as Figure 18-10 shows, protostars take many thousands or millions of years to evolve significantly. Instead, we must compare different star clusters at various stages in their evolution to piece together the history of star formation in a cluster.

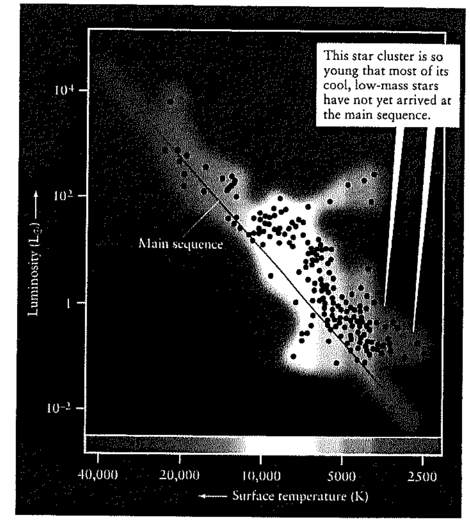
All the stars in a cluster may begin to form nearly simultaneously, but they do not all become main-sequence stars at the same time. As you can see from their evolutionary tracks (see Figure 18-10), high-mass stars evolve more rapidly than low-mass stars. The more massive the protostar, the sooner it develops the central pressures and temperatures needed for steady hydrogen fusion to begin, thus joining the main sequence.

Upon reaching the main sequence, *high-mass* protostars become hot, ultraluminous stars of spectral types O and B. As we saw in Section 18-2, these types of stars have ultraviolet radiation that ionize the surrounding interstellar medium to produce an H II region. Figure 18-17 shows such an H II region, called the Eagle Nebula, surrounding the young star cluster M16. A few hundred thousand years ago, this region of space would have had a far less dramatic appearance. It was then a dark nebula, with protostars just beginning to form. Over the intervening millennia, mass ejection from these evolving protostars swept away the obscuring dust. The exposed young, hot stars heated the relatively thin remnants of the original dark nebula, creating the H II region that we see today.

When the most massive protostars to form out of a dark nebula have reached the main sequence, other *low-mass* protostars are still evolving nearby within their dusty cocoons. The evolution of these low-mass stars can be disturbed by their more massive neighbors. As an example, the inset in Figure 18-17 is a close-up of part of the Eagle Nebula. Within these opaque pillars of cold gas and dust, protostars are still forming. At the same time, however, the pillars are being eroded by intense ultraviolet light from hot, massive stars that have already shed their cocoons.



(a) The star cluster NGC 2264 R I **V** U X G




(b) An H-R diagram of the stars in NGC 2264

**Figure 18-18**

A Young Star Cluster and Its H-R Diagram **(a)** This photograph shows an H II region and the young star cluster NGC 2264 in the constellation Monoceros (the Unicorn). It lies about 800 pc (2600 ly) from Earth. **(b)** Each dot plotted on this H-R diagram represents a star in NGC 2264 whose luminosity and surface temperature have been determined. This star cluster probably started forming only 2 million years ago. (Anglo-Australian Observatory)

As each pillar evaporates, the embryonic stars within have their surrounding material stripped away prematurely, limiting the total mass that these stars can accrete.

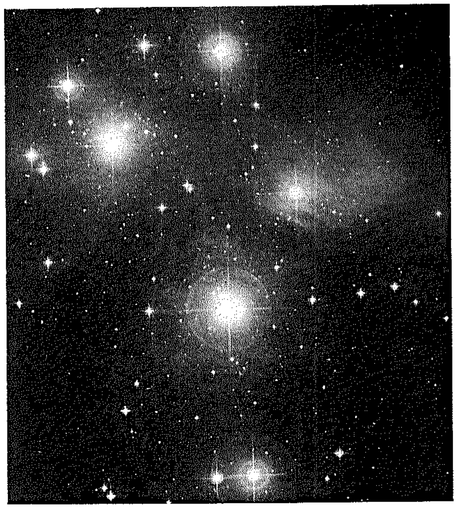
### Analyzing Young Clusters Using H-R Diagrams

Star clusters tell us still more about how high-mass and low-mass stars evolve. Figure 18-18*a* shows the young star cluster NGC 2264 and its associated emission nebula. Astronomers have measured each star's apparent brightness and color ratio. Knowing the distance to the cluster, they have deduced the luminosities and surface temperatures of the stars (see Section 17-2 and Section 17-4). Figure 18-18*b* shows all these stars on an H-R diagram. Note that the hottest and most massive stars, with surface temperatures around 20,000 K, are on the main sequence. Stars cooler than about 10,000 K, however, have not yet quite arrived at the main sequence. These are less massive stars in the final stages of pre–main-sequence contraction and are just now beginning to ignite thermonuclear reactions at their centers. To find the ages of these stars, we can compare Figure 18-18*b* with the theoretical calculations of protostar evolution in Figure 18-10. It turns out that this particular cluster is probably about 2 million years old.

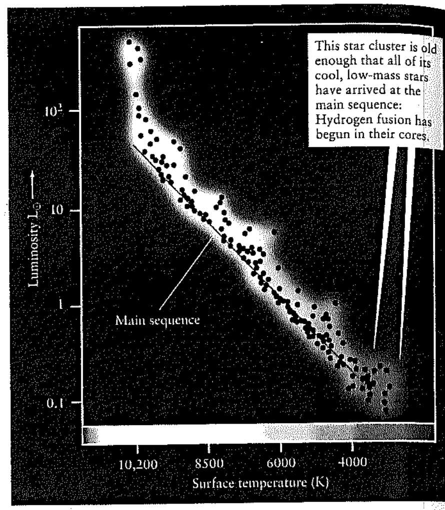
**The H-R diagram of a young cluster reveals how much time has elapsed since its stars began to form**

Figure 18-19*a* shows another young star cluster called the Pleiades. The photograph shows gas that must once have formed an H II region around this cluster and has dissipated into interstellar space, leaving only traces of dusty material that forms reflection nebulae around the cluster's stars. This implies that the Pleiades must be older than NGC 2264, the cluster in Figure 18-18*a*, which is still surrounded by an H II region. The H-R diagram for the Pleiades in Figure 18-19*b* bears out this idea. In contrast to the H-R diagram for NGC 2264, nearly all the stars in the Pleiades are on the main sequence. The cluster's age is about 50 million years, which is how long it takes for the least massive stars to finally begin hydrogen fusion in their cores.

**CAUTION!** Note that the data points for the most massive stars in the Pleiades (at the upper left of the H-R diagram in Figure



(a) The Pleiades star cluster R I V U X G




(b) An H-R diagram of the stars in the Pleiades

## Figure 18-19

The Pleiades and Its H-R Diagram (a) The Pleiades star cluster is 117 pc (380 ly) from Earth in the constellation Taurus, and can be seen with the naked eye. (b) Each dot plotted on this H-R diagram represents a star in the Pleiades whose luminosity and surface temperature have been measured. (*Note:* The scales on this H-R diagram are different from those in Figure 18-18*b*.) The Pleiades is about 50 million ($5 \times 10^7$) years old. (Anglo-Australian Observatory)

18-19*b*) lie above the main sequence. This is *not* because these stars have yet to arrive at the main sequence. Rather, these stars were the first members of the cluster to arrive at the main sequence some time ago and are now the first members to leave it. They have used up the hydrogen in their cores, so the steady process of core hydrogen fusion that characterizes main-sequence stars cannot continue. In Chapter 19 we will see why massive stars spend a rather short time as main-sequence stars and will study what happens to stars after the main-sequence phase of their lives.

A loose collection of stars such as NGC 2264 or the Pleiades is referred to as an **open cluster** (or *galactic cluster,* since such clusters are usually found in the plane of the Milky Way Galaxy). Open clusters possess barely enough mass to hold themselves together by gravitation. Occasionally, a star moving faster than average will escape, or "evaporate," from an open cluster. Indeed, by the time the stars are a few billion years old, they may be so widely separated that a cluster no longer exists.

If a group of stars is gravitationally unbound from the very beginning—that is, if the stars are moving away from one another so rapidly that gravitational forces cannot keep them together—then the group is called a **stellar association.** Because young stellar associations are typically dominated by luminous O and B main-sequence stars, they are also called **OB associations.** The image that opens this chapter shows part of an OB association in the southern constellation Ara (the Altar).

## 18-7 Star birth can begin in giant molecular clouds

We have seen that star formation takes place within dark nebulae. But where within our Galaxy are these dark nebulae found? Does star formation take place everywhere within the Milky Way, or only in certain special locations? The answers to such questions can enhance our understanding of star formation and of the nature of our home Galaxy.

### Exploring the Interstellar Medium at Millimeter Wavelengths

Dark nebulae are a challenge to locate simply because they are dark—they do not emit visible light. Nearby dark nebulae can be seen silhouetted against background stars or H II regions (see Figure 18-2), but sufficiently distant dark nebulae are impossible to see in contrast with background visible light because of interstellar extinction from dust grains. They can, however, be detected using longer-wavelength radiation that can pass unaffected through interstellar dust. In fact, dark nebulae actually emit radiation at millimeter wavelengths.

Such emission takes place because in the cold depths of interstellar space, atoms combine to form molecules. The laws of quantum mechanics predict that just as electrons within atoms can occupy only certain specific energy levels (see Section 5-8), molecules can vibrate and rotate only at certain specific rates. When a molecule goes from one vibrational state or rotational state to another, it either emits or absorbs a photon. (In the same way, an atom emits or absorbs a photon as an electron jumps from one energy level to another.) Most molecules are strong emitters of radiation with wavelengths of around 1 to 10 millimeters (mm). Consequently, observations with radio telescopes tuned to millimeter wavelengths make it possible to detect interstellar molecules of different types. More than 100 different kinds of molecules have so far been discovered in interstellar space, and the list is constantly growing.

**Observing the Galaxy at millimeter wavelengths reveals the cold gas that spawns new stars**

Hydrogen is by far the most abundant element in the universe. Unfortunately, in cold nebulae much of it is in a molecular form ($H_2$) that is difficult to detect. The reason is that the hydrogen molecule is symmetric, with two atoms of equal mass joined together, and such molecules do not emit many photons at radio frequencies. In contrast, asymmetric molecules that consist of two atoms of unequal mass joined together, such as carbon monoxide (CO), are easily detectable at radio frequencies. When a carbon monoxide molecule makes a transition from one rate of rotation to another, it emits a photon at a wavelength of 2.6 mm or shorter.

The ratio of carbon monoxide to hydrogen in interstellar space is reasonably constant: For every CO molecule, there are about 10,000 $H_2$ molecules. As a result, carbon monoxide is an excellent "tracer" for molecular hydrogen gas. Wherever astronomers detect strong emission from CO, they know molecular hydrogen gas must be abundant.

### Giant Molecular Clouds

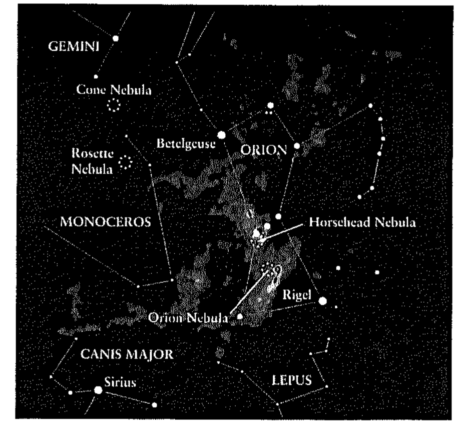
The first systematic surveys of our Galaxy looking for 2.6-mm CO radiation were undertaken in 1974 by the American astronomers Philip Solomon and Nicholas Scoville. In mapping the locations of CO emission, they discovered huge clouds, now called **giant molecular clouds,** that must contain enormous amounts of hydrogen. These clouds have masses in the range of $10^5$ to $2 \times 10^6$ solar masses and diameters that range from about 15 to 100 pc (50 to 300 ly). Inside one of these clouds, there are about 200 hydrogen molecules per cubic centimeter. This density is several thousand times greater than the average density of matter in the disk of our Galaxy, yet only $10^{-17}$ as dense as the air we breathe. Astronomers now estimate that our Galaxy contains about 5000 of these enormous clouds.

Figure 18-20 is a map of radio emissions from carbon monoxide in the constellations Orion and Monoceros. Note the extensive areas of the sky covered by giant molecular clouds. This part of the sky is of particular interest because it includes several star-forming regions. By comparing the radio map with the star chart overlay, you can see that the areas where CO emission is strongest, and, thus, where giant molecular clouds are densest, are sites of star formation. Therefore, giant molecular clouds are associated with the formation of stars. Particularly dense regions within these clouds form dark nebulae, and within these stars are born.

By using CO emissions to map out giant molecular clouds, astronomers can find the locations in our Galaxy where star formation occurs. These investigations reveal that molecular clouds clearly outline our Galaxy's spiral arms, as Figure 18-21 shows. These clouds lie roughly 1000 pc (3000 ly) apart and are strung along the spiral arms like beads on a string. This arrangement resembles the spacing of H II regions along the arms of other spiral galaxies, such as the galaxy shown in Figure 18-8*a*. The presence of both molecular clouds and H II regions shows that spiral arms are sites of ongoing star formation.

### Star Formation in Spiral Arms

In Chapter 23 we will learn that spiral arms are locations where matter "piles up" temporarily as it orbits the center of the Galaxy.




## Figure 18-20 R I V U X G

Mapping Molecular Clouds A radio telescope was tuned to a wavelength of 2.6 mm to detect emissions from carbon monoxide (CO) molecules in the constellations Orion and Monoceros. The result was this false-color map, which shows a 35° × 40° section of the sky. The Orion and Horsehead star-forming nebulae are located at sites of intense CO emission (shown in red and yellow), indicating the presence of a particularly dense molecular cloud at these sites of star formation. The molecular cloud is much thinner at the positions of the Cone and Rosette nebulae, where star formation is less intense. (Courtesy of R. Maddalena, M. Morris, J. Moscowitz, and P. Thaddeus) for observing with a telescope? Explain how you determined this.

54. Use the *Starry Night Enthusiast*™ program to examine the Milky Way Galaxy. Open the **Favorites** pane and click on **Stars > Sun in Milky Way** to display our Galaxy from a position 0.150 million light-years above the galactic plane. (You can remove the astronaut's feet from this view if desired by clicking on **View > Feet.**) You can zoom in or out on the Galaxy using the + and − buttons at the upper right end of the toolbar. You can move the Galaxy by holding down the mouse button while moving the mouse. You can also rotate the Galaxy by putting the mouse cursor over the image and holding down the **Shift** key while holding down the mouse button and moving the mouse. (a) You can identify H II regions by their characteristic magenta color. Describe where in the Galaxy you find these. Are most found in the inner part of the Galaxy or in its outer regions? (b) Where do you find dark lanes of dust—in the inner part of the Galaxy or in its outer regions? Do you see any connection between the locations of dust and of H II regions? If there is a connection, what do you think causes it? If there is not a connection, why is this the case? You can examine the location of our Galaxy in relation to neighboring galaxies by turning the Milky Way edge-on and by increasing the distance from Earth using the **up** key below the Viewing Location on the toolbar.

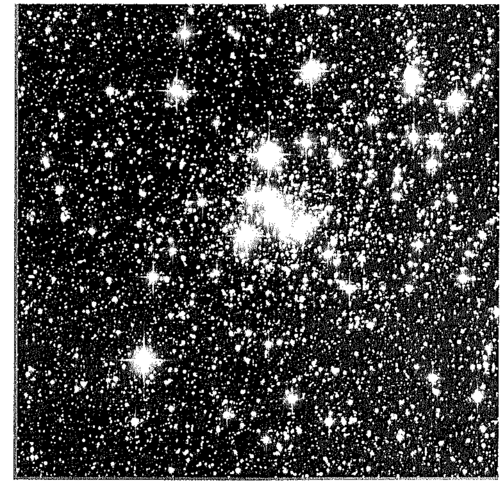
### Collaborative Exercises

55. Imagine that your group walks into a store that specializes in selling antique clothing. Prepare a list of observable characteristics that you would look for to distinguish which items were from the early, middle, and late twentieth century. Also, write a paragraph that specifically describes how this task is similar to how astronomers understand the evolution of stars.

56. Consider advertisement signs visible at night in your community and provide specific examples of ones that are examples of the three different types of nebulae that astronomers observe and study. If an example doesn't exist in your community, creatively design an advertisement sign that could serve as an example.

57. The pre–main-sequence evolutionary tracks shown in Figure 18-10 describe the tracks of seven protostars of different masses. Imagine a new sort of H-R diagram that plots a human male's increasing age versus decreasing hair density on the head instead of increasing luminosity versus decreasing temperature. Create and carefully label a sketch of this imaginary HaiR diagram showing both the majority of the U.S. male population and a few oddities. Finally, draw a line that clearly labels your sketch to show how a typical male undergoing male-pattern baldness might slowly change position on the HaiR diagram over the course of a human life span.

# 19 Stellar Evolution: On and After the Main Sequence



R I V U X G

The red stars in this image of open cluster NGC 290 are red giants, a late stage in stellar evolution. (ESA/NASA/Edward W. Olszewski, U. of Arizona)

Imagine a world like Earth, but orbiting a star more than 100 times larger and 2000 times more luminous than our Sun. Bathed in the star's intense light, the surface of this world is utterly dry, airless, and hot enough to melt iron. If you could somehow survive on the daytime side of this world, you would see the star filling almost the entire sky.

This bizarre planet is not a creation of science fiction—it is our own Earth some 7.6 billion years from now. The bloated star is our own Sun, which in that remote era will have become a red-giant star.

In this chapter, we'll learn how a main-sequence star evolves into a red giant when all the hydrogen in its core is consumed. The star's core contracts and heats up, but its outer layers expand and cool. In the hot, compressed core, helium fusion becomes a new energy source. The more massive a star, the more rapidly it consumes its core's hydrogen and the sooner it evolves into a giant.

The interiors of stars are hidden from our direct view, so much of the story in this chapter is based on theory. We back up those calculations with observations of star clusters, which contain stars of different masses but roughly the same age. (An example is the cluster shown here, many of whose stars have evolved into luminous red giants.) Other observations show that some red-giant stars actually pulsate, and that stars can evolve along very different paths if they are part of a binary star system.

## 19-1 During a star's main-sequence lifetime, it expands and becomes more luminous

In their cores, main-sequence stars are all fundamentally alike. As we saw in Section 18-4, it is in their cores that all such stars convert hydrogen into helium by thermonuclear reactions. This process is called **core hydrogen fusion**. The total time that a star will spend fusing hydrogen into helium in its core, and thus the total time that it will spend as a main-sequence star, is called its **main-sequence lifetime**. For our Sun, the main-sequence lifetime is about 12 billion ($1.2 \times 10^{10}$) years. Hydrogen

**Over the past 4.56 billion years, thermonuclear reactions have caused an accumulation of helium in our Sun's core**

## Learning Goals

*By reading the sections of this chapter, you will learn*

19-1 How a main-sequence star changes as it converts hydrogen to helium

19-2 What happens to a star when it runs out of hydrogen fuel

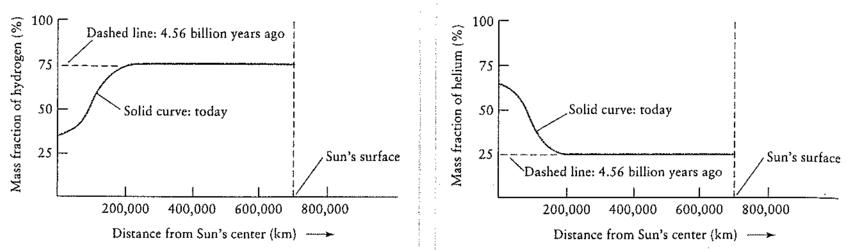
19-3 How aging stars can initiate a second stage of thermonuclear fusion

19-4 How H-R diagrams for star clusters reveal the later stages in the evolution of stars

19-5 The two kinds of stellar populations and their significance

19-6 Why some aging stars pulsate and vary in luminosity

19-7 How stars in a binary system can evolve very differently from single, isolated stars




(a) Hydrogen in the Sun's interior

(b) Helium in the Sun's interior

**Figure 19-1**

Changes in the Sun's Chemical Composition These graphs show the percentage by mass of **(a)** hydrogen and **(b)** helium at different points within the Sun's interior. The dashed horizontal lines show that these percentages were the same throughout the Sun's volume when it first formed. As the solid curves show, over the past $4.56 \times 10^9$ years, thermonuclear reactions at the core have depleted hydrogen in the core and increased the amount of helium in the core.

fusion has been going on in the Sun's core for the past 4.56 billion ($4.56 \times 10^9$) years, so our Sun is less than halfway through its main-sequence lifetime.

What happens to a star like the Sun after the core hydrogen has been used up, so that it is no longer a main-sequence star? As we will see, it expands dramatically to become a red giant. To understand why this happens, it is useful to first look at how a star evolves *during* its main-sequence lifetime. The nature of that evolution depends on whether its mass is less than or greater than about 0.4 $M_\odot$.

## Main-Sequence Stars of 0.4 $M_\odot$ or Greater: Consuming Core Hydrogen

A protostar becomes a main-sequence star when steady hydrogen fusion begins in its core and it achieves *hydrostatic equilibrium*—a balance between the inward force of gravity and the outward pressure produced by hydrogen fusion (see Section 16-2 and Section 18-4). Such a freshly formed main-sequence star is called a **zero-age main-sequence star.**

We make the distinction between "main sequence" and "zero-age main sequence" because a star undergoes noticeable changes in luminosity, surface temperature, and radius during its main-sequence lifetime. These changes are a result of core hydrogen fusion, which alters the chemical composition of the core. As an example, when our Sun first formed, its composition was the same at all points throughout its volume: by mass, about 74% hydrogen, 25% helium, and 1% heavy elements. But as Figure 19-1 shows, the Sun's core now contains a greater mass of helium than of hydrogen. (There is still enough hydrogen in the Sun's core for another 7 billion years or so of core hydrogen fusion.)

**CAUTION!** Although the outer layers of the Sun are also predominantly hydrogen, there are two reasons why this hydrogen cannot undergo fusion. The first reason is that while the temperature and pressure in the core are high enough for thermonuclear reactions to take place, the temperatures and pressure in the outer layers are not. The second reason is that there is no flow of material between the Sun's core and outer layers, so the hydrogen in the outer layers cannot move into the hot, high-pressure core to undergo fusion. The same is true for main-sequence stars with masses of about 0.4 $M_\odot$ or greater. (We will see below that the outer layers *can* undergo fusion in main-sequence stars with a mass less than about 0.4 $M_\odot$.)

Thanks to core hydrogen fusion, the total number of atomic nuclei in a star's core decreases with time: In each reaction, four hydrogen nuclei are converted to a single helium nucleus (see the *Cosmic Connections* figure in Section 16-1, as well as Box 16-1). With fewer particles bouncing around to provide the core's internal pressure, the core contracts slightly under the weight of the star's outer layers. Compression makes the core denser and increases its temperature. (Box 19-1 gives some everyday examples of how the temperature of a gas changes when it compresses or expands.) As a result of these changes in density and temperature, the pressure in the compressed core is actually higher than before.

As the star's core shrinks, its outer layers expand and shine more brightly. Here's why: As the core's density and temperature increase, hydrogen nuclei in the core collide with one another more frequently, and the rate of core hydrogen fusion increases. Hence, the star's luminosity increases. The radius of the star as a whole also increases slightly, because increased core pressure pushes outward on the star's outer layers. The star's surface temperature changes as well, because it is related to the luminosity and radius (see Section 17-6 and Box 17-4). As an example, theoretical calculations indicate that over the past $4.56 \times 10^9$ years, our Sun has become 40% more luminous, grown in radius by 6%, and increased in surface temperature by 300 K (Figure 19-2).

As a main-sequence star ages and evolves, the increase in energy outflow from its core also heats the material immediately surrounding the core. As a result, hydrogen fusion can begin in this surrounding material. By tapping this fresh supply of hydrogen, a star manages to eke out a few million more years of main-sequence existence.



**BOX 19-1** Astronomy Down to Earth

### Compressing and Expanding Gases

As a star evolves, various parts of the star either contract or expand. When this happens, the gases behave in much the same way as gases here on Earth when they are forced to compress or allowed to expand.

When a gas is compressed, its temperature rises. You know this by personal experience if you have ever had to inflate a bicycle tire with a hand pump. As you pump, the compressed air gets warm and makes the pump warm to the touch. The same effect happens on a larger scale in southern California during Santa Ana winds or downwind from the Rocky Mountains when there are Chinook winds. Both of these strong winds blow from the mountains down to the lowlands. Even though the mountain air is cold, the winds that reach low elevations can be very hot. (Chinook winds have been known to raise the temperature by as much as 27°C, or 49°F, in only 2 minutes!) The explanation is compression. Air blown downhill by the winds is compressed by the greater air pressure at lower altitudes, and this compression raises the temperature of the air.

Expanding gases tend to drop in temperature. When you open a bottle of carbonated beverage, the gases trapped in the bottle expand and cool down. The cooling can be so great that a little cloud forms within the neck of the bottle. Clouds form in the atmosphere in the same way. Rising air cools as it goes to higher altitudes, where the pressure is lower, and the cooling makes water in the air condense into droplets.

Here's an experiment you can do to feel the cooling of expanding gases. Your breath is actually quite warm, as you can feel if you open your mouth wide, hold the back of your hand next to your mouth, and exhale. But if you bring your lips together to form an "o" and again blow on your hand, your breath feels cool. In the second case, your exhaled breath has to expand as it passes between your lips to the outside, which makes its temperature drop.

## Main-Sequence Stars of Less than 0.4 $M_\odot$: Consuming All Their Hydrogen

The story is somewhat different for the least massive main-sequence stars, with masses between 0.08 $M_\odot$ (the minimum mass for sustained thermonuclear reactions to take place in a star's core) and about 0.4 $M_\odot$. These stars, of spectral class M, are called **red dwarfs** because they are small in size and have a red color due to their low surface temperature. They are also very numerous; about 85% of all stars in the Milky Way Galaxy are red dwarfs.

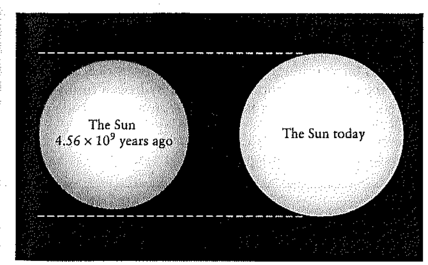



**Figure 19-2**

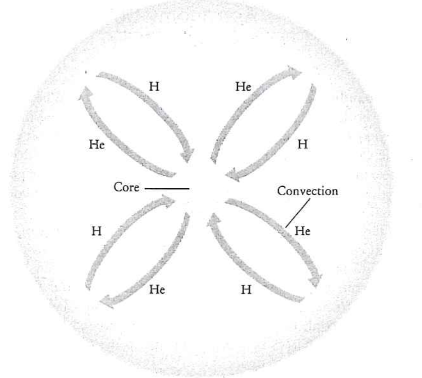
The Zero-Age Sun and Today's Sun Over the past $4.56 \times 10^9$ years, much of the hydrogen in the Sun's core has been converted into helium, the core has contracted a bit, and the Sun's luminosity has gone up by about 40%. These changes in the core have made the Sun's outer layers expand in radius by 6% and increased the surface temperature from 5500 K to 5800 K.

In a red dwarf, helium does *not* accumulate in the core to the same extent as in the Sun's core. The reason is that in a red dwarf there are convection cells of rising and falling gas that extend throughout the star's volume and penetrate into the core (see Figure 18-12*c*). These convection cells drag helium outward from the core and replace it with hydrogen from the outer layers (Figure 19-3). The fresh hydrogen can undergo thermonuclear fusion that releases energy and makes additional helium. This helium is then dragged out of the core by convection and replaced by even more hydrogen from the red dwarf's outer layers.

As a consequence, over a red dwarf's main-sequence lifetime essentially all of the star's hydrogen can be consumed and converted to helium. The core temperature and pressure in a red dwarf is less than in the Sun, so thermonuclear reactions happen more slowly than in our Sun. Calculations indicate that it takes hundreds of billions of years for a red dwarf to convert all of its hydrogen completely to helium. The present age of the universe is only 13.7 billion years, so there has not yet been time for any red dwarfs to become pure helium.

## A Star's Mass Determines Its Main-Sequence Lifetime

The main-sequence lifetime of a star depends critically on its mass. As Table 19-1 shows, massive stars have short main-sequence lifetimes because they are also very luminous (see Section 17-9, and particularly the *Cosmic Connections* figure for Chapter 17). To emit energy so rapidly, these stars must be




## Figure 19-3

A Fully Convective Red Dwarf In a red dwarf—a main-sequence star with less than about 0.4 solar masses—helium (He) created in the core by thermonuclear reactions is carried to the star's outer layers by convection. Convection also brings fresh hydrogen (H) from the outer layers into the core. This process continues until the entire star is helium.

depleting the hydrogen in their cores at a prodigious rate. Hence, even though a massive O or B main-sequence star contains much more hydrogen fuel in its core than is in the entire volume of a red dwarf of spectral class M, the O or B star exhausts its hydrogen much sooner. High-mass O and B stars gobble up the available hydrogen fuel in only a few million years, while red dwarf stars of very low mass take hundreds of billions of years to use up their hydrogen. Thus, a main-sequence star's mass determines not only its luminosity and spectral type, but also how long it can remain a main-sequence star (see Box 19-2 for details).

We saw in Section 18-4 how more-massive stars evolve more quickly through the protostar phase to become main-sequence stars (see Figure 18-10). In general, the more massive the star, the more rapidly it goes through *all* the phases of its life. Still, most of the stars we are able to detect are in their main-sequence phase, because this phase lasts so much longer than other luminous phases.

In the remainder of this chapter we will look at the luminous phases that can take place after the end of a star's main-sequence lifetime. (In Chapters 20, 21, and 22 we will explore the final phases of a star's existence, when it ceases to have an appreciable luminosity.)

## 19-2 When core hydrogen fusion ceases, a main-sequence star like the Sun becomes a red giant

TUTORIAL 19.1

Like so many properties of stars, what happens at the end of a star's main-sequence lifetime depends on its mass. If the star is a red dwarf of less than about 0.4 $M_\odot$, after hundreds of billions of years the star has converted all of its hydrogen to helium. It is possible for helium to undergo thermonuclear fusion, but this requires temperatures and pressures far higher than those found within a red dwarf. Thus, this red dwarf will end its life as an inert ball of helium, which has no further nuclear reactions, but still glows due to its internal heat. As it radiates energy into space, it slowly cools and shrinks. This slow, quiet demise is the ultimate fate of the 85% of stars in the Milky Way that are red dwarfs. (As we have seen, there has not yet been time in the history of the universe for any red dwarf to reach this final stage in its evolution.)

What is the fate of stars more massive than about 0.4 $M_\odot$, including the Sun? As we will see, the late stages of their evolution are far more dramatic. Studying these stages will give us insight into the fate of our solar system and of life on Earth.

### Stars of 0.4 $M_\odot$ or Greater: From Main-Sequence Star to Red Giant

When a star of at least 0.4 solar masses reaches the end of its main-sequence lifetime, all of the hydrogen in its core has been used up and hydrogen fusion ceases there. In this new stage, hydrogen fusion continues only in the hydrogen-rich material just outside the core, a situation called **shell hydrogen fusion.** At first, this process occurs only in the hottest region just outside the core, where the hydrogen fuel has not yet been exhausted. Outside this region, no fusion reactions take place.

Strangely enough, the end of the core hydrogen fusion process leads to an *increase* in the core's temperature. Here's why: When thermonuclear reactions first cease in the core, nothing remains to generate heat there. Hence, the core starts to cool and the pressure in the core starts to decrease. This pressure decrease allows the star's core to again compress under the weight of the outer layers. As the core contracts, its temperature again increases, and heat begins to flow outward from the core even though no nuclear reactions are taking place there. (Technically, gravitational energy is converted into thermal energy, as in Kelvin-Helmholtz contraction; see Section 8-4 and Section 16-1).

**In the transition from main-sequence star to red giant, the star's core contracts while its outer layers expand**

This new flow of heat warms the gases around the core, increasing the rate of shell hydrogen fusion and making the shell eat further outward into the surrounding matter. Helium produced by reactions in the shell falls down onto the core, which continues

### Table 19-1 Approximate Main-Sequence Lifetimes

| Mass ($M_\odot$) | Surface temperature (K) | Spectral class | Luminosity ($L_\odot$) | Main-sequence lifetime ($10^6$ years) |
|---|---|---|---|---|
| 25 | 35,000 | O | 80,000 | 4 |
| 15 | 30,000 | B | 10,000 | 15 |
| 3 | 11,000 | A | 60 | 800 |
| 1.5 | 7000 | F | 5 | 4500 |
| 1.0 | 6000 | G | 1 | 12,000 |
| 0.75 | 5000 | K | 0.5 | 25,000 |
| 0.50 | 4000 | M | 0.03 | 700,000 |

*The main-sequence lifetimes were estimated using the relationship $t \propto 1/M^{2.5}$ (see Box 19-2).*



BOX 19-2 Tools of the Astronomer's Trade

## Main-Sequence Lifetimes

Hydrogen fusion converts a portion of a star's mass into energy. We can use Einstein's famous equation relating mass and energy to calculate how long a star will remain on the main sequence.

Suppose that $M$ is the mass of a star and $f$ is the fraction of the star's mass that is converted into energy by hydrogen fusion. During its main-sequence lifetime, the total energy $E$ supplied by the hydrogen fusion can be expressed as

$$E = fMc^2$$

In this equation $c$ is the speed of light.

This energy from hydrogen fusion is released gradually over millions or billions of years. If $L$ is the star's luminosity (energy released per unit time) and $t$ is the star's main-sequence lifetime (the total time over which the hydrogen fusion occurs), then we can write

$$L = \frac{E}{t}$$

(Actually, this equation is only an approximation. A star's luminosity is not quite constant over its entire main-sequence lifetime. But the variations are not important for our purposes.) We can rewrite this equation as

$$E = Lt$$

From this equation and $E = fMc^2$, we see that

$$Lt = fMc^2$$

We can rearrange this equation as

$$t = \frac{fMc^2}{L}$$

Thus, a star's lifetime on the main sequence is proportional to its mass ($M$) divided by its luminosity ($L$). Using the symbol $\propto$ to denote "is proportional to," we write

$$t \propto \frac{M}{L}$$

We can carry this analysis further by recalling that main-sequence stars obey the mass-luminosity relation (see Section 17-9, especially the *Cosmic Connections* figure). The distribution of data on the graph in the *Cosmic Connections* figure in Section 17-9 tells us that a star's luminosity is roughly proportional to the 3.5 power of its mass:

$$L \propto M^{3.5}$$

Substituting this relationship into the previous proportionality, we find that

$$t \propto \frac{M}{M^{3.5}} = \frac{1}{M^{2.5}} = \frac{1}{M^2\sqrt{M}}$$

This approximate relationship can be used to obtain rough estimates of how long a star will remain on the main sequence. It is often convenient to relate these estimates to the Sun (a typical 1-$M_\odot$ star), which will spend $1.2 \times 10^{10}$ years on the main sequence.

EXAMPLE: How long will a star whose mass is 4 $M_\odot$ remain on the main sequence?

**Situation:** Given the mass of a star, we are asked to determine its main-sequence lifetime.

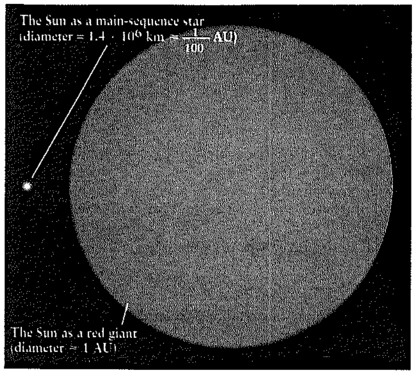
**Tools:** We use the relationship $t \propto 1/M^{2.5}$.

**Answer:** The star has 4 times the mass of the Sun, so it will be on the main sequence for approximately

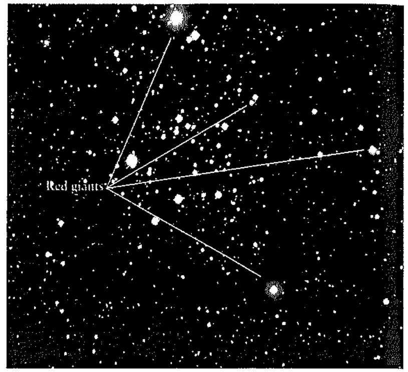
$$\frac{1}{4^{2.5}} = \frac{1}{4^2\sqrt{4}} = \frac{1}{32} \text{ times the Sun's main-sequence lifetime}$$

Thus, a 4-$M_\odot$ main-sequence star will fuse hydrogen in its core for about $(1/32) \times 1.2 \times 10^{10}$ years, or about $4 \times 10^8$ (400 million) years.

**Review:** Our result makes sense: A star more massive than the Sun must have a shorter main-sequence lifetime.




(a) The Sun today and as a red giant




(b) Red giant stars in the star cluster M50 R I V U X G

**Figure 19-4**

Red Giants (a) The present-day Sun produces energy in a hydrogen-fusing core about 100,000 km in diameter. Some 7.6 billion years from now, when the Sun becomes a red giant, its energy source will be a shell only about 30,000 km in diameter within which hydrogen fusion will take place at a furious rate. The Sun's luminosity will be about 2000 times greater than today, and the increased luminosity will make the Sun's outer layers expand to approximately 100 times their present size. (b) This composite of visible and infrared images shows bright red giant stars in the open cluster M50 in the constellation Monoceros (the Unicorn). (T. Credner and S. Kohle, Calar Alto Observatory)

to contract and heat up as it gains mass. Over the course of hundreds of millions of years, the core of a 1-$M_\odot$ star compresses to about one-third of its original radius, while its central temperature increases from about 15 million ($1.5 \times 10^7$) K to about 100 million ($10^8$) K.

During this post–main-sequence phase, the star's outer layers expand just as dramatically as the core contracts. As the hydrogen-fusing shell works its way outward, egged on by heat from the contracting core, the star's luminosity increases substantially. This increases the star's internal pressure and makes the star's outer layers expand to many times the original radius. This tremendous expansion causes those outer layers to cool down, and the star's surface temperature drops (see Box 19-1). Once the temperature of the star's bloated surface falls to about 3500 K, the gases glow with a reddish hue, in accordance with Wien's law (see Figure 17-7*a*). The star is then appropriately called a red giant (Figure 19-4). Thus, we see that red-giant stars are former main-sequence stars that have evolved into a new stage of existence. We can summarize these observations as a general rule:

***Stars join the main sequence when they begin hydrogen fusion in their cores. They leave the main sequence and become giant stars when the core hydrogen is depleted.***

Red-giant stars undergo substantial **mass loss** because of their large diameters and correspondingly weak surface gravity. This makes it relatively easy for gases to escape from the red giant into space. Mass loss can be detected in a star's spectrum, because gas escaping from a red giant toward a telescope on Earth produces narrow absorption lines that are slightly blueshifted by the Doppler effect (review Figure 5-26). Typical observed blueshifts correspond to a speed of about 10 km/s. A typical red giant loses roughly $10^{-7}$ $M_\odot$ of matter per year. For comparison, the Sun's present-day mass loss rate is only $10^{-14}$ $M_\odot$ per year. Hence, an evolving star loses a substantial amount of mass as it becomes a red giant. Figure 19-5 shows a star losing mass in this way.

## The Distant Future of Our Solar System

We can use these ideas to peer into the future of our planet and our solar system. The Sun's luminosity will continue to increase as it goes through its main-sequence lifetime, and the temperature of Earth will increase with it. One and a half billion years from now Earth's average surface temperature will be 50°C (122°F). By $3\frac{1}{2}$ billion years from now the surface temperature of Earth will exceed the boiling temperature of water. All the oceans will boil away, and Earth will become utterly incapable of supporting life. These increasingly hostile conditions will pose the ultimate challenge to whatever intelligent beings might inhabit Earth in the distant future.

About 7 billion years from now, our Sun will finish converting hydrogen into helium at its core. As the Sun's core contracts, its atmosphere will expand to envelop Mercury and perhaps reach to the orbit of Venus. Roughly 700 million years after leaving the main sequence, our red-giant Sun will have swollen to a diameter of about 1 AU—roughly 100 times larger than its present size—and its surface temperature will have dropped to about 3500 K. Shell hydrogen fusion will proceed at such a furious rate that our star will shine with the brightness of 2000 present-day Suns. Some of the inner planets will be vaporized, and the thick atmospheres of the outer planets will evaporate away to reveal tiny, rocky cores. Thus, in its later years, the aging Sun may destroy the planets that have accompanied it since its birth.

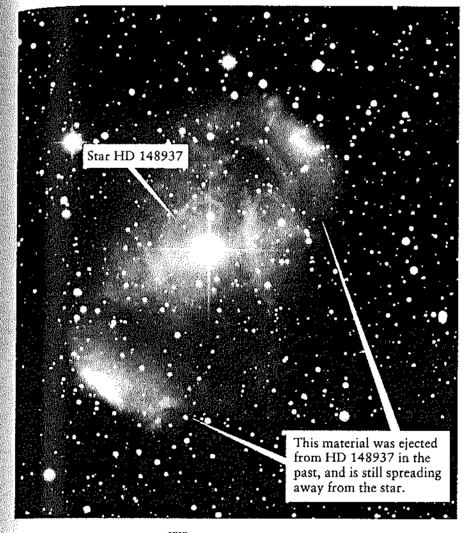



**Figure 19-5** R I V U X G

A Mass-Loss Star As stars age and become giant stars, they expand tremendously and shed matter into space. This star, HD 148937, is losing matter at a high rate. Other strong outbursts in the past ejected the clouds that surround HD 148937. These clouds absorb ultraviolet radiation from the star, which excites the atoms in the clouds and causes them to glow. The characteristic red color of the clouds reveals the presence of hydrogen (see Section 5-6) that was ejected from the star's outer layers. (David Malin, Anglo-Australian Observatory)

## 19-3 Fusion of helium into carbon and oxygen begins at the center of a red giant

When a star with a mass greater than 0.4 $M_\odot$ first changes from a main-sequence star (Figure 19-6*a*) to a red giant (Figure 19-6*b*), its hydrogen-fusing shell surrounds a small, compact core of almost pure helium. In a red giant of moderately low mass, which the Sun will become 7 billion years from now, the dense helium core is about twice the diameter of Earth. Most of this core helium was *produced* by thermonuclear reactions during the star's main-sequence lifetime; during the red-giant era, this helium will *undergo* thermonuclear reactions.

### Core Helium Fusion

Helium, the "ash" of hydrogen fusion, is a potential nuclear fuel: **Helium fusion,** the thermonuclear fusion of helium nuclei to make heavier nuclei, releases energy. But this reaction cannot take place within the core of our present-day Sun because the temperature there is too low. Each helium nucleus contains two protons, so it has twice the positive electric charge of a hydrogen nucleus, and there is a much stronger electric repulsion between two helium nuclei than between two hydrogen nuclei. For helium nuclei to overcome this repulsion and get close enough to fuse together, they must be moving at very high speeds, which means that the temperature of the helium gas must be very high. (For more on the relationship between the temperature of a gas and the speed of atoms in the gas, see Box 7-2.)

When a star first becomes a red giant, the temperature of the contracted helium core is still too low for helium nuclei to fuse. But as the hydrogen-fusing shell adds mass to the helium core, the core contracts even more, further increasing the star's central temperature. When the central temperature finally reaches 100 million ($10^8$) K, **core helium fusion**—that is, thermonuclear fusion of helium in the core—begins. As a result, the aging star again has a central energy source for the first time since leaving the main sequence (Figure 19-6*c*).

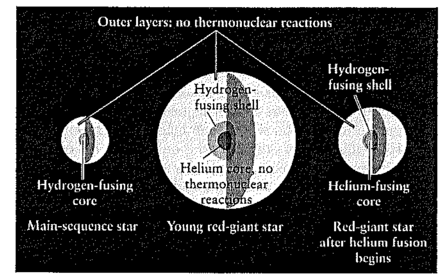



**Figure 19-6**

Stages in the Evolution of a Star with More than 0.4 Solar Masses (a) During the star's main-sequence lifetime, hydrogen is converted into helium in the star's core. (b) When the core hydrogen is exhausted, hydrogen fusion continues in a shell, and the star expands to become a red giant. (c) When the temperature in the red giant's core becomes high enough because of contraction, core helium fusion begins (right). (These three pictures are *not* drawn to scale. The star is about 100 times larger in its red-giant phase than in its main-sequence phase, then shrinks somewhat when core helium fusion begins.)

Helium fusion occurs in two steps. First, two helium nuclei combine to form a beryllium nucleus:

$$^{4}He + ^{4}He \rightarrow ^{8}Be$$

This particular beryllium isotope, which has four protons and four neutrons, is very unstable and breaks into two helium nuclei soon after it forms. However, in the star's dense core a third helium nucleus may strike the $^{8}Be$ nucleus before it has a chance to fall apart. Such a collision creates a stable isotope of carbon and releases energy in the form of a gamma-ray photon ($\gamma$):

$$^{8}Be + ^{4}He \rightarrow ^{12}C + \gamma$$

This process of fusing three helium nuclei to form a carbon nucleus is called the **triple alpha process**, because helium nuclei ($^{4}He$) are also called **alpha particles** by nuclear physicists. Some of the carbon nuclei created in this process can fuse with an additional helium nucleus to produce a stable isotope of oxygen and release more energy:

$$^{12}C + ^{4}He \rightarrow ^{16}O + \gamma$$

Thus, both carbon and oxygen make up the "ash" of helium fusion. The *Cosmic Connections* figure summarizes the reactions involved in helium fusion.

It is interesting to note that $^{12}C$ and $^{16}O$ are the most common isotopes of carbon and oxygen, respectively; the vast majority of the carbon atoms in your body are $^{12}C$, and almost all the oxygen you breathe is $^{16}O$. We will discuss the significance of this in Section 19-5.

The second step in the triple alpha process and the process of oxygen formation both release energy. The onset of these reactions reestablishes thermal equilibrium and prevents any further gravitational contraction of the star's core. A mature red giant fuses helium in its core for a much shorter time than it spent fusing hydrogen in its core as a main-sequence star. For example, in the distant future the Sun will sustain helium core fusion for only about 100 million years—this period is only about 1% of the time that hydrogen fusion occurs. (While this is going on, hydrogen fusion is still continuing in a shell around the core.)

## The Helium Flash and Electron Degeneracy

How helium fusion begins at a red giant's center depends on the mass of the star. In high-mass red giants (greater than about 2 to 3 $M_{\odot}$), helium fusion begins gradually as temperatures in the star's core approach $10^8$ K. In red giants with a mass less than about 2 to 3 $M_{\odot}$, helium fusion begins explosively and suddenly, in what is called the **helium flash.** Table 19-2 summarizes these differences.

The helium flash occurs because of unusual conditions that develop in the core of a moderately low-mass star as it becomes a red giant. To appreciate these conditions we must first understand how an ordinary gas behaves. Then we can explore how the densely packed electrons at the star's center alter this behavior.

When a gas is compressed, it usually becomes denser and warmer. To describe this process, scientists use the convenient concept of an **ideal gas,** which has a simple relationship between pressure, temperature, and density. Specifically, the pressure exerted by an ideal gas is directly proportional to both the density and the temperature of the gas. Many real gases actually behave like an ideal gas over a wide range of temperatures and densities.

**Table 19-2 How Helium Core Fusion Begins in Different Red Giants**

| Mass of star | Onset of helium fusion in core |
|---|---|
| More than about 0.4 but less than 2–3 solar masses | Explosive (helium flash) |
| More than 2–3 solar masses | Gradual |

*Stars with less than about 0.4 solar masses do not become red giants (see Section 19-2).*

Under most circumstances, the gases inside a star act like an ideal gas. If the gas expands, it cools down, and if it is compressed, it heats up (see Box 19-1). This behavior serves as a safety valve, ensuring that the star remains in thermal equilibrium (see Section 16-2). For example, if the rate of thermonuclear reactions in the star's core should increase, the additional energy releases heat and expands the core. This expansion cools the core's gases and slows the rate of thermonuclear reactions back to the original value. Conversely, if the rate of thermonuclear reactions should decrease, the core will cool down and compress under the pressure of the overlying layers. The compression of the core will make its temperature increase, thus speeding up the thermonuclear reactions and returning them to their original rate.

In a red giant with a mass between about 0.4 $M_{\odot}$ and 2-3 $M_{\odot}$, however, the core behaves very differently from an ideal gas. The core must be compressed tremendously in order to become hot enough for helium fusion to begin. At these extreme pressures and temperatures, the atoms are completely ionized, and most of the core consists of nuclei and detached electrons. Eventually, the free electrons become so closely crowded that a limit to further compression is reached, as predicted by a remarkable law of quantum mechanics called the **Pauli exclusion principle.** Formulated in 1925 by the Austrian physicist Wolfgang Pauli, this principle states that two electrons cannot simultaneously occupy the same quantum state. A quantum state is a particular set of circumstances concerning locations and speeds that are available to a particle. In the submicroscopic world of electrons, the Pauli exclusion principle is analogous to saying that you can't have two things in the same place at the same time.

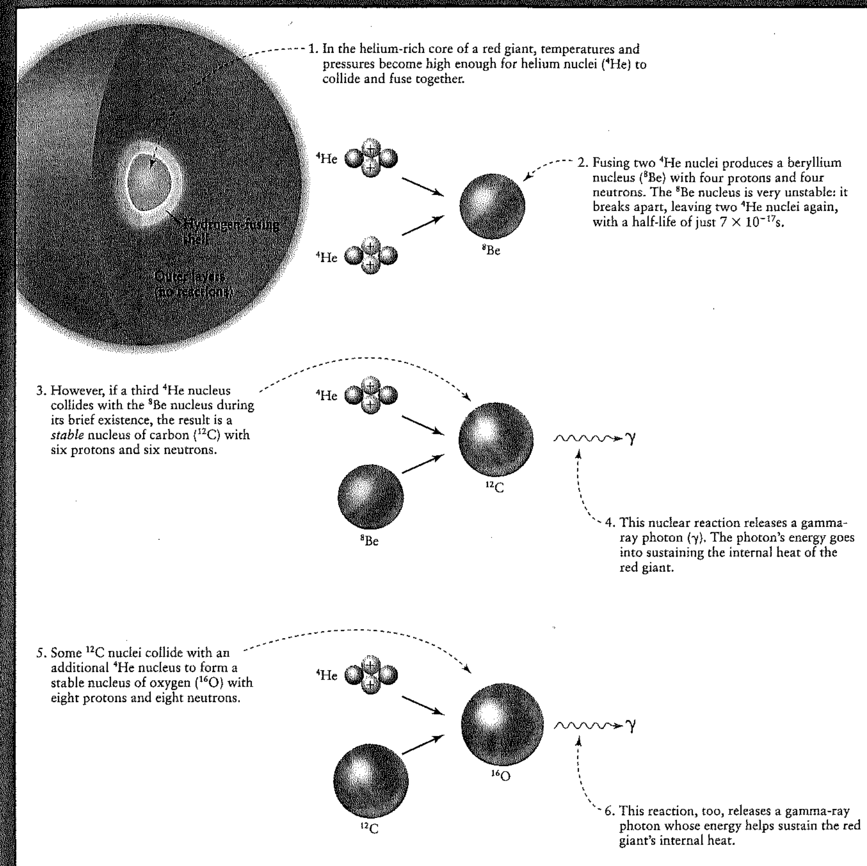
**Before helium fusion begins in a red giant's core, the helium gas there behaves like a metal**

Just before the onset of helium fusion, the electrons in the core of a low-mass star are so closely crowded together that any further compression would violate the Pauli exclusion principle. Because the electrons cannot be squeezed any closer together, they produce a powerful pressure that resists further core contraction.

This phenomenon, in which closely packed particles resist compression as a consequence of the Pauli exclusion principle, is called **degeneracy.** Astronomers say that the electrons in the helium-rich core of a low-mass red giant are "degenerate," and

# COSMIC CONNECTIONS Helium Fusion In A Red Giant

A star becomes a red giant after the fusion of hydrogen into helium in its core has come to an end. As the red giant's core shrinks and heats up, a new cycle of reactions can occur that create the even heavier elements carbon and oxygen.



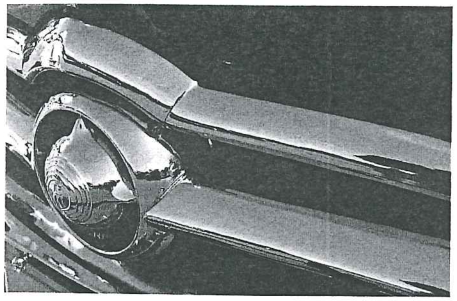


Figure 19-7 R I V U X G

Degenerate Electrons The electrons in an ordinary piece of metal, like the chrome grille on this classic car, are so close together that they are affected by the Pauli exclusion principle. The resulting degenerate electron pressure helps make metals strong and difficult to compress. A more powerful version of this same effect happens inside the cores of low-mass red-giant stars. (Santokh Kochar/PhotoDisc)

that the core is supported by **degenerate-electron pressure.** This degenerate pressure, unlike the pressure of an ideal gas, does not depend on temperature. Remarkably, you can find degenerate electrons on Earth in an ordinary piece of metal (Figure 19-7).

When the temperature in the core of a low-mass red giant reaches the high level required for the triple alpha process, energy begins to be released. The helium heats up, which makes the triple alpha process happen even faster. However, the pressure provided by the degenerate electrons is independent of the temperature, so the pressure does not change. Without the "safety valve" of increasing pressure, the star's core cannot expand and cool. The rising temperature causes the helium to fuse at an ever-increasing rate, producing the helium flash.

Eventually, the temperature becomes so high that the electrons in the core are no longer degenerate. The electrons then behave like an ideal gas and the star's core expands, terminating the helium flash. These events occur so rapidly that the helium flash is over in seconds, after which the star's core settles down to a steady rate of helium fusion.

**CAUTION!** The term "helium flash" might give you the impression that a star emits a sudden flash of light when the helium flash occurs. If this were true, it would be an incredible sight. During the brief time interval when the helium flash occurs, the helium-fusing core is $10^{11}$ times more luminous than the present-day Sun, which is similar to the total luminosity of all the stars in the Milky Way Galaxy! But, in fact, the helium flash has no immediately visible consequences—for two reasons. First, much of the energy released during the helium flash goes into heating the core and terminating the degenerate state of the electrons. Second, the energy that does escape the core is largely absorbed by the star's outer layers, which are quite opaque (just like the Sun's present-day interior; see Section 16-2). Therefore, the explosive drama of the helium flash takes place where it cannot be seen directly.

## The Continuing Evolution of a Red Giant

Whether a helium flash occurs or not, the onset of core helium fusion actually causes a *decrease* in the luminosity of the star. This decrease is the opposite of what you might expect—after all, turning on a new energy source should make the luminosity greater, not less. What happens is that after the onset of core helium fusion, a star's superheated core expands like an ideal gas. (If the star is of sufficiently low mass to have had a degenerate core, the increased temperature after the helium flash makes the core too hot to remain degenerate. Hence, these stars also end up with cores that behave like ideal gases.) Temperatures drop around the expanding core, so the hydrogen-fusing shell reduces its energy output and the star's luminosity decreases. This temperature decrease allows the star's outer layers to contract and heat up. Consequently, a post–helium-flash star is less luminous, hotter at the surface, and smaller than a red giant.

Core helium fusion lasts for only a relatively short time. Calculations suggest that a $1\text{-}M_\odot$ star like the Sun sustains core hydrogen fusion for about 12 billion ($1.2 \times 10^{10}$) years, followed by about 250 million ($2.5 \times 10^{8}$) years of shell hydrogen fusion leading up to the helium flash. After the helium flash, such a star can fuse helium in its core (while simultaneously fusing hydrogen in a shell around the core) for only 100 million ($10^{8}$) years, a mere 1% of its main-sequence lifetime. Figure 19-8 summarizes these evolutionary stages in the life of a $1\text{-}M_\odot$ star. In Chapter 20 we will take up the story of what happens after a star has consumed all the helium in its core.

Here is the story of post–main-sequence evolution in its briefest form: Before the beginning of core helium fusion, the star's core compresses and the outer layers expand, and just after core helium fusion begins, the core expands and the outer layers compress. We will see in Chapter 20 that this behavior, in which the inner and outer regions of the star change in opposite ways, occurs again and again in the final stages of a star's evolution.

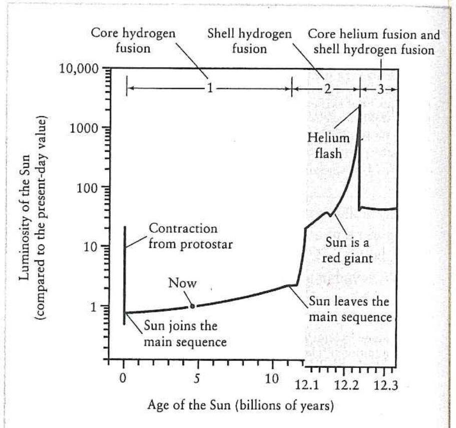



Figure 19-8

Stages in the Evolution of the Sun This diagram shows how the luminosity of the Sun (a $1\text{-}M_\odot$ star) changes over time. The Sun began as a protostar whose luminosity decreased rapidly as the protostar contracted. Once established as a main-sequence star with core hydrogen fusion, the Sun's luminosity increases slowly over billions of years. The post-main-sequence evolution is much more rapid, so a different time scale is used in the right-hand portion of the graph. (Adapted from Mark A. Garlick, based on calculations by I.-Juliana Sackmann and Kathleen E. Kramer)



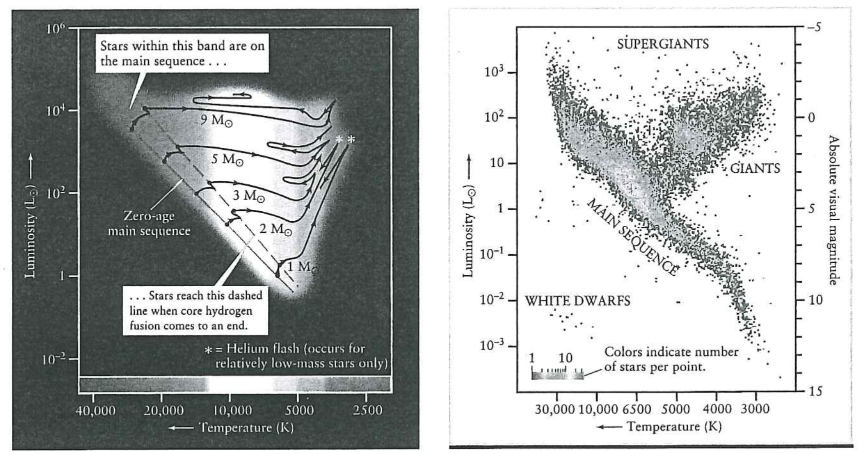
## 19-4 H-R diagrams and observations of star clusters reveal how red giants evolve

TUTORIAL 19.2

To see how stars evolve during and after their main-sequence lifetimes, it is helpful to follow them on a Hertzsprung-Russell (H-R) diagram. On such a diagram, zero-age main-sequence stars lie along a line called the **zero-age main sequence,** or **ZAMS** (Figure 19-9*a*). These stars have just emerged from their protostar stage, are steadily fusing hydrogen into helium in their cores, and have attained hydrostatic equilibrium. With the passage of time, hydrogen in a main-sequence star's core is converted to helium, the luminosity slowly increases, the star slowly expands, and the star's position on the H-R diagram inches away from the ZAMS. As a result, the main sequence on an H-R diagram is a fairly broad band rather than a narrow line (Figure 19-9*b*).

## Post-Main-Sequence Evolution on an H-R Diagram

The dashed line in Figure 19-9*a* denotes stars whose cores have been exhausted of hydrogen and in which core hydrogen fusion has ceased. These stars have reached the ends of their main-sequence lifetimes. From there, the points representing high-mass stars ($3\ M_\odot$, $5\ M_\odot$, and $9\ M_\odot$) move rapidly from left to right across the H-R diagram. This means that, although the star's surface temperature is decreasing, its surface area is increasing at a rate that keeps its overall luminosity roughly constant. During this transition, the star's core contracts and its outer layers expand as energy flows outward from the hydrogen-fusing shell.




(a) Post-main-sequence evolutionary tracks of five stars with different mass

(b) H-R diagram of 20,853 stars—note the width of the main sequence

WEB LINK 19.3

Figure 19-9

H-R Diagrams of Stellar Evolution on and off the Main Sequence **(a)** The two lowest-mass stars shown here ($1\ M_\odot$ and $2\ M_\odot$) undergo a helium flash at their centers, as shown by the asterisks. In the high-mass stars, core helium fusion ignites more gradually where the evolutionary tracks make a sharp downward turn in the red-giant region on the right hand side of the H-R diagram. **(b)** Data from the *Hipparcos* satellite (see Section 17-1) was used to create this H-R diagram. The thickness of the main sequence is due in large part to stars evolving during their main-sequence lifetimes. (a: Adapted from I. Iben; b: Adapted from M. A. C. Perryman)

Just before core helium fusion begins, the evolutionary tracks of high-mass stars turn upward in the red-giant region of the H-R diagram (to the upper right of the main sequence). After core helium fusion begins, however, the cores of these stars expand, the outer layers contract, and the evolutionary tracks back away from these temporary peak luminosities. The tracks then wander back and forth in the red-giant region while the stars readjust to their new energy sources.

Figure 19-9*a* also shows the evolutionary tracks of two stars of moderately low mass (1 $M_\odot$ and 2 $M_\odot$). The onset of core helium fusion in these stars occurs with a helium flash, indicated by the red asterisks in the figure. As we saw in the previous section, after the helium flash, these stars shrink and become less luminous. The decrease in size is proportionately greater than the decrease in luminosity, and so the surface temperatures increase. Hence, after the helium flash, the evolutionary tracks for the 1-$M_\odot$ and 2-$M_\odot$ stars move down and to the left.

## A Simulated Star Cluster: Tracking $4\frac{1}{2}$ Billion Years of Stellar Evolution

We can summarize our understanding of stellar evolution from birth through the onset of helium fusion by following the evolution of a hypothetical cluster of stars. We saw in Section 18-6 that the stars that make up a cluster all begin to form at essentially the same time but have different initial masses. Hence, studying star clusters allows us to compare how stars of different masses evolve.

The eight H-R diagrams in Figure 19-10 are from a computer simulation of the evolution of 100 stars that all form at the same moment and differ only in initial mass. All 100 stars begin as cool protostars on the right side of the H-R diagram (see Figure 19-10*a*). The protostars are spread out on the diagram according to their masses, and the greater the mass, the greater the protostar's initial luminosity. As we saw in Section 18-3, the source of a protostar's luminosity is its gravitational energy. As the protostar contracts, this gravitational energy is converted to thermal energy and radiated into space.

The most massive protostars contract and heat up very rapidly. After only 5000 years, they have already moved across the H-R diagram toward the main sequence (see Figure 19-10*b*). After 100,000 years, these massive stars have ignited hydrogen fusion in their cores and have settled down on the main sequence as O stars (see Figure 19-10*c*). After 3 million years, stars of moderate mass have also ignited core hydrogen fusion and become main-sequence stars of spectral classes B and A (see Figure 19-10*d*). Meanwhile, low-mass protostars continue to inch their way toward the main sequence as they leisurely contract and heat up.

After 30 million years (see Figure 19-10*e*), the most massive stars have depleted the hydrogen in their cores and become red giants. These stars have moved from the upper left end of the main sequence to the upper right corner of the H-R diagram. (This simulation follows stars only to the red-giant stage, after which they are simply deleted from the diagram.) Intermediate-mass stars lie on the main sequence, while the lowest-mass stars are still in the protostar stage and lie above the main sequence.

After 66 million years (see Figure 19-10*f*), even the lowest-mass protostars have finally ignited core hydrogen fusion and have settled down on the main sequence as cool, dim, M stars. These lowest-mass stars can continue to fuse hydrogen in their cores for hundreds of billions of years.

In the final two H-R diagrams (Figures 19-10*g* and 19-10*h*), the main sequence stars get "peeled" or "eaten away" from the upper left to the lower right as stars exhaust their core supplies of hydrogen and evolve into red giants. The stars that leave the main sequence between Figure 19-10*g* and Figure 19-10*h* have masses between about 1 $M_\odot$ and 3 $M_\odot$ and undergo the helium flash in their cores.

For all stars in this simulation, the giant stage lasts only a brief time compared to the star's main-sequence lifetime. Compared to a 1-$M_\odot$ star (see Figure 19-8), a more massive star has a shorter main-sequence lifetime *and* spends a shorter time as a giant star. Thus, at any given time, only a small fraction of the stellar population is passing through the giant stage. Hence, most of the stars we can see through telescopes are main-sequence stars. As an example, of the stars within 4.00 pc (13.05 ly) of the Sun listed in Appendix 4, only one—Procyon A—is presently evolving from a main-sequence star into a giant. Two other nearby stars are white dwarfs, an even later stage in stellar evolution that we will discuss in Chapter 20. (By contrast, most of the *brightest* stars listed in Appendix 5 are giants and supergiants. Although they make up only a small fraction of the stellar population, these stars stand out due to their extreme luminosity.)

## Real Star Clusters: Cluster Ages and Turnoff Points

WEB LINK 19.4

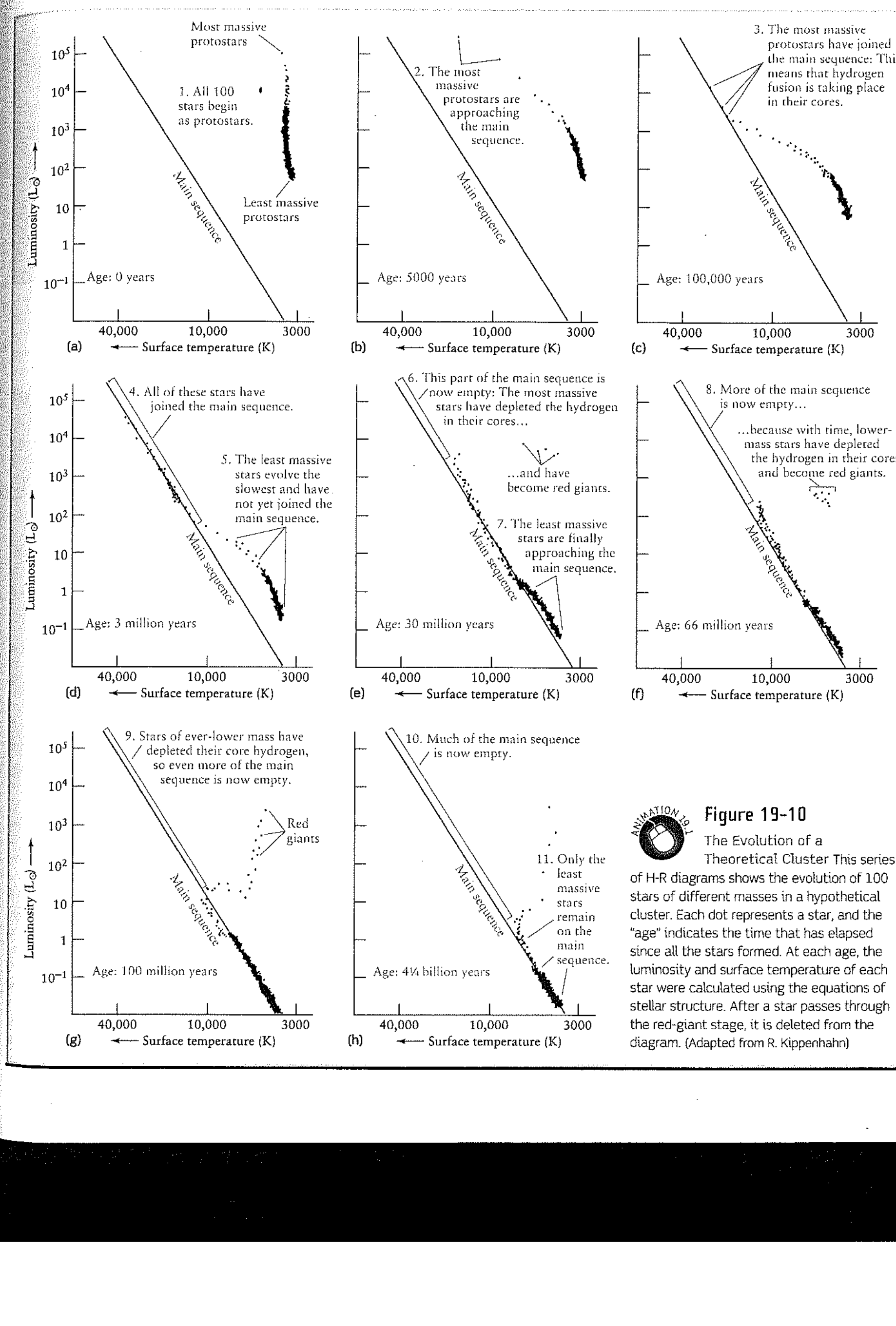
The evolution of the hypothetical cluster displayed in Figure 19-10 helps us interpret what we see in actual star clusters. We can observe the early stages of stellar evolution in **open clusters**, which typically contain a few hundred to a few thousand stars. Many open clusters are just a few million years old, so their H-R diagrams resemble Figure 19-10*d*, 19-10*e*, or 19-10*f* (see Section 18-6, especially Figure 18-18 and Figure 18-19).

Figure 19-11 shows two open clusters of different ages. The nearer cluster, called M35, must be relatively young because it contains several dozen luminous, blue, high-mass main-sequence stars. These stars lie in the upper part of the main sequence on an H-R diagram. They have main-sequence lifetimes of only a few hundred million years, so M35 can be no older than that. Some of the most luminous stars in M35 are red or yellow in color; these are stars that ended their main-sequence lifetimes some time ago and have evolved into red giants. The H-R diagram for this cluster resembles Figure 19-10*g*.

**As a cluster ages and its stars end their main-sequence lifetimes, the cluster's color changes from blue to red**

There are no high-mass blue main-sequence stars at all in NGC 2158, the more distant cluster shown in Figure 19-11. Any such stars that were once in NGC 2158 have long since come to the end of their main-sequence lifetimes. As a result, the main sequence in this cluster has been "eaten away" more than that of M35, leaving only stars that are yellow or red in color. This tells us that NGC 2158 must be older than M35 (compare Figures






ANIMATION 19.1

**Figure 19-10**

The Evolution of a Theoretical Cluster This series of H-R diagrams shows the evolution of 100 stars of different masses in a hypothetical cluster. Each dot represents a star, and the "age" indicates the time that has elapsed since all the stars formed. At each age, the luminosity and surface temperature of each star were calculated using the equations of stellar structure. After a star passes through the red-giant stage, it is deleted from the diagram. (Adapted from R. Kippenhahn)

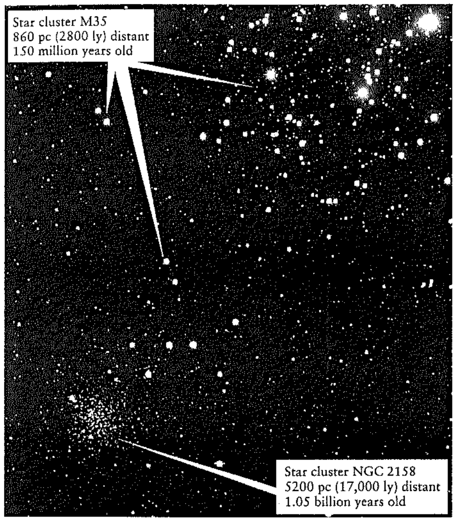



**Figure 19-11** R I V U X G
Two Open Clusters The two clusters in this image, M35 and NGC 2158, lie in almost the same direction in the constellation Gemini. The nearer cluster, M35, has a number of luminous blue main-sequence stars with surface temperatures around 10,000 K as well as a few red giants. Hence its H-R diagram resembles that shown in Figure 19-10*g*, and its age is around 100 million years (more accurately, around 150 million). The more distant cluster, NGC 2158, has *no* blue main-sequence stars; long ago, all of these massive stars came to the end of their main-sequence lifetimes and became giants. The H-R diagram for NGC 2158 is intermediate between Figure 19-10*g* and Figure 19-10*h*, and its age (1.05 billion years) is as well. (Canada-France-Hawaii Telescope; J.-C. Cuillandre, CFHT; and Coelum)

19-10*g* and 19-10*h*). This example shows that as a cluster ages, it generally becomes redder in its average color.

We can see even later stages in stellar evolution by studying **globular clusters,** so called because of their spherical shape. A typical globular cluster contains up to 1 million stars in a volume less than 100 parsecs across (Figure 19-12). Among these are many highly evolved post–main-sequence stars.

Globular clusters must be old, because they contain no high-mass main-sequence stars. To determine that these clusters are old, you would measure the apparent magnitude (a measure of apparent brightness, which we introduced in Section 17-3) and color ratio of many stars in a globular cluster, then plot the data as shown in Figure 19-13. Such a **color-magnitude diagram** for a cluster is equivalent to an H-R diagram. The color ratio of a star tells you its surface temperature (as described in Section 17-4), and because all the stars in the cluster are at essentially the same distance from us, their relative brightnesses indicate their relative luminosities. What you would discover is that a globular cluster's main sequence has been "peeled" or "eaten away" even more extensively than the open cluster NGC 2158 in Figure 19-11. Hence, globular clusters must be even older than NGC 2158. In a typical globular cluster, all the main-sequence stars with masses more than about 1M$_\odot$ or 2 M$_\odot$ evolved long ago into red giants. Only low-mass, slowly evolving stars still have core hydrogen fusion. (Compare Figure 19-13 with Figure 19-10*h*.)

**CAUTION!** The inset in Figure 19-12 shows something surprising: There are luminous *blue* stars in the ancient globular cluster M10. This seems to contradict our earlier statements that blue main-sequence stars evolve into red giants after just a few hundred million years. The explanation is that these are *not* main-sequence stars, but rather **horizontal-branch stars.** These stars get their name because in the color-magnitude diagram of a globular cluster, they form a horizontal grouping in the left-of-center portion of the diagram (see Figure 19-13). Horizontal-branch stars are relatively low-mass stars that have already become red giants and undergone a helium flash, so there is both core helium fusion and shell hydrogen fusion taking place in their interiors. After the helium flash their luminosity decreased to about 50 L$_\odot$ (compared to about 1000 L$_\odot$ before the flash) and their outer layers contracted and heated, giving these stars their blue color. In years to come, these stars will move back toward the red-giant region as their fuel is devoured. Our own Sun will go through a horizontal-branch phase in the distant future; this is the phase labeled by the number 3 at the far right of Figure 19-8.

## Measuring the Ages of Star Clusters

The idea that a cluster's main sequence is progressively "eaten away" is the key to determining the age of a cluster. In the H-R diagram for a very young cluster, all the stars are on or near the main sequence. (An example is the open cluster NGC 2264, shown in Figure 18-18.) As a cluster gets older, however, stars begin to leave the main sequence. The high-mass, high-luminosity stars are the first to consume their core hydrogen and become red giants. As time passes the main sequence gets shorter and shorter, like a candle burning down (see parts *d* through *h* of Figure 19-10).

The age of a cluster can be found from the **turnoff point,** which is the top of the surviving portion of the main sequence stars on the cluster's H-R diagram (see Figure 19-13). The stars at the turnoff point are just now exhausting the hydrogen in their cores, so their main-sequence lifetime is equal to the age of the cluster. For example, in the case of the globular cluster M55 plotted in Figure 19-13, 0.8-M$_\odot$ stars have just left the main sequence, indicating that the cluster's age is more than 12 billion ($1.2 \times 10^{10}$) years (see Table 19-1).

Figure 19-14 shows data for several star clusters plotted on a single H-R diagram. This graph also shows turnoff-point times from which the ages of the clusters can be estimated.



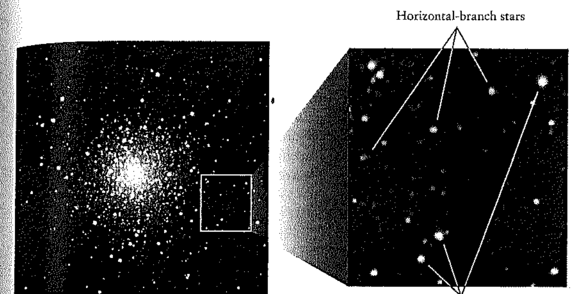



**Figure 19-12** R I V U X G
A Globular Cluster This cluster, called M10, contains a few hundred thousand stars within a diameter of only 20 pc (70 ly). It lies approximately 5000 pc (16,000 ly) from Earth in the constellation Ophiuchus (the Serpent Holder). Most of the stars in this image are either red giants or blue, horizontal-branch stars with both core helium fusion and shell hydrogen fusion. (T. Credner and S. Kohle, Astronomical Institutes of the University of Bonn)

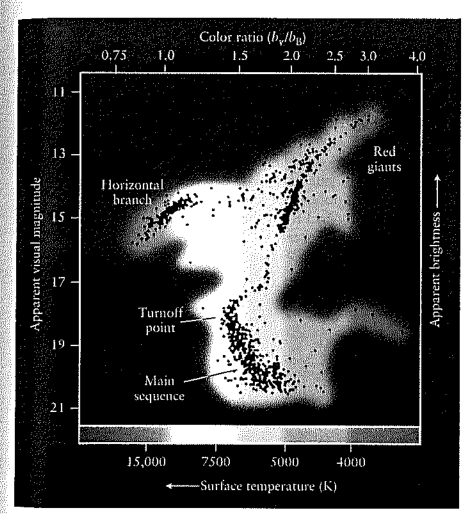



**Figure 19-13**
A Color-Magnitude Diagram of a Globular Cluster Each dot in this diagram represents the apparent visual magnitude (a measure of the brightness as seen through a V filter) and surface temperature (as measured by the color ratio $b_V/b_B$) of a star in the globular cluster M55 in Sagittarius. Because all the stars in M55 are at essentially the same distance from Earth (about 6000 pc or 20,000 ly), their apparent visual magnitudes are a direct measure of their luminosities. Note that the upper half of the main sequence is missing. (Adapted from D. Schade, D. VandenBerg, and F. Hartwick)

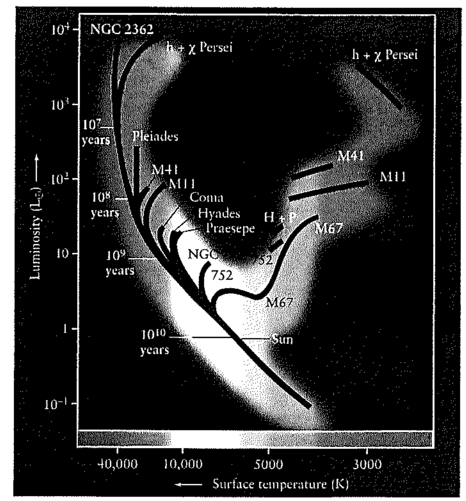



**Figure 19-14**
An H-R Diagram for Open Star Clusters The black bands indicate where stars from various open clusters fall on the H-R diagram. The age of a cluster can be estimated from the location of the cluster's turnoff point, where the cluster's most massive stars are just now leaving the main sequence. The times for these turnoff points are listed alongside the main sequence. For example, the Pleiades cluster turnoff point is near the $10^8$-year point, so this cluster is about $10^8$ years old. (Adapted from A. Sandage)

## 19-5 Stellar evolution has produced two distinct populations of stars

Studies of star clusters reveal a curious difference between the youngest and oldest stars in our Galaxy. Stars in the youngest clusters (those with most of their main sequences still intact) are said to be **metal rich**, because their spectra contain many prominent spectral lines of heavy elements. (Recall from Section 17-5 that astronomers use the term "metal" to denote any element other than hydrogen and helium, which are the two lightest elements.) Such stars are also called **Population I stars.** The Sun is a relatively young, metal-rich, Population I star.

By contrast, the spectra of stars in the oldest clusters show only weak lines of heavy elements. These ancient stars are thus said to be **metal poor,** because heavy elements are only about 3% as abundant in these stars as in the Sun. They are also called **Population II stars.** The stars in globular clusters are metal poor, Population II stars. Figure 19-15 shows the difference in spectra between a metal-poor, Population II star and the Sun (a metal-rich, Population I star).

**CAUTION!** Note that "metal rich" and "metal poor" are relative terms. In even the most metal-rich star known, metals make up just a few percent of the total mass of the star.

### Stellar Populations and the Origin of Heavy Elements

To explain why there are two distinct populations of stars, we must go back to the Big Bang, the explosive origin of the universe that took place some 13.7 billion years ago. As we will discuss in Chapter 27, the early universe consisted almost exclusively of hydrogen and helium, with almost no heavy elements (metals). The first stars to form were likewise metal poor. The least massive of these stars have survived to the present day and are now the ancient stars of Population II.

The more massive of the original stars evolved more rapidly and no longer shine. But as these stars evolved, helium fusion in their cores produced metals—carbon and oxygen. In the most massive stars, as we will learn in Chapter 21, further thermonuclear reactions produced even heavier elements. As these massive original stars aged and died, they expelled their metal-enriched gases into space. (The star shown in Figure 19-5 is going through such a mass-loss phase late in its life.) This expelled material joined the interstellar medium and was eventually incorporated into a second generation of stars that have a higher concentration of heavy elements. These metal-rich members of the second stellar generation are the Population I stars, of which our Sun is an example.

**Stars like the Sun contain material that was processed through an earlier generation of stars**

**CAUTION!** Be careful not to let the designations of the two stellar populations confuse you. Population *I* stars are members of a *second* stellar generation, while Population *II* stars belong to an older *first* generation.

The relatively high concentration of heavy elements in the Sun means that the solar nebula, from which both the Sun and planets formed (see Section 8-4), must likewise have been metal rich. Earth is composed almost entirely of heavy elements, as are our bodies. Thus, our very existence is intimately linked to the Sun's being a Population I star. A planet like Earth probably could not have formed from the metal-poor gases that went into making Population II stars.

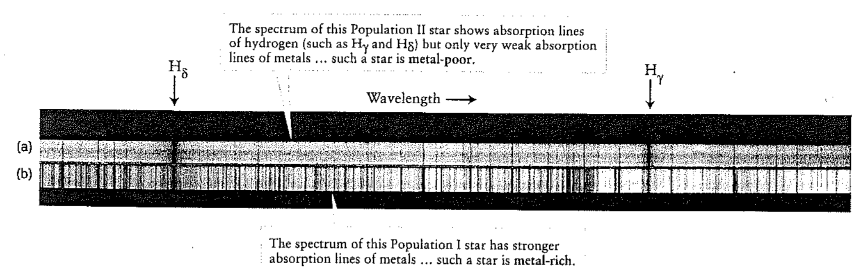



**Figure 19-15** R I V U X G

Spectra of a Metal-Poor Star and a Metal-Rich Star The abundance of metals (elements heavier than hydrogen and helium) in a star can be inferred from its spectrum. These spectra compare **(a)** a metal-poor, Population II star and **(b)** a metal-rich, Population I star (the Sun) of the same surface temperature. We described the hydrogen absorption lines $H_\gamma$ (wavelength 434 nm) and $H_\delta$ (wavelength 410 nm) in Section 5-8. (Lick Observatory)



The concept of two stellar populations provides insight into our own origins. Recall from Section 19-3 that helium fusion in red-giant stars produces the same isotopes of carbon ($^{12}C$) and oxygen ($^{16}O$) that are found most commonly on Earth. The reason is that Earth's carbon and oxygen atoms, including all of those in your body, actually *were* produced by helium fusion. These reactions occurred billions of years ago within an earlier generation of stars that died and gave up their atoms to the interstellar medium—the same atoms that later became part of our solar system, our planet, and our bodies. We are literally children of the stars.

## 19-6 Many mature stars pulsate

ANIMATION 19.2

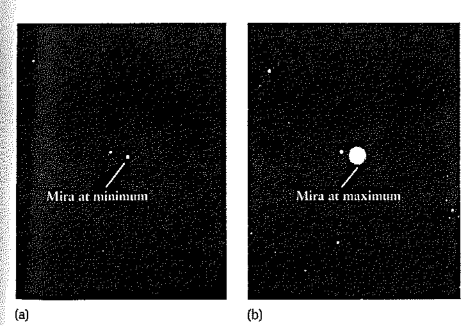
We saw in Section 16-3 that the surface of our Sun vibrates in and out, although by only a small amount. But other stars undergo substantial changes in size, alternately swelling and shrinking. As these stars pulsate, they also vary dramatically in brightness. We now understand that these **pulsating variable stars** are actually evolved, post–main-sequence stars.

### Long-Period Variables

Pulsating variable stars were first discovered in 1595 by David Fabricius, a Dutch minister and amateur astronomer. He noticed that the star o (omicron) Ceti is sometimes bright enough to be easily seen with the naked eye but at other times fades to invisibility (Figure 19-16). By 1660, astronomers realized that these brightness variations repeated with a period of 332 days. Seventeenth-century astronomers were so enthralled by this variable star that they renamed it Mira ("wonderful").

Mira is an example of a class of pulsating stars called **long-period variables.** These stars are cool red giants that vary in brightness by a factor of 100 or more over a period of months or years. With surface temperatures of about 3500 K and average luminosities that range from about 10 to 10,000 $L_\odot$, they occupy the upper right side of the H-R diagram (Figure 19-17). Some, like Mira, are periodic, but others are irregular. Many eject large amounts of gas and dust into space.

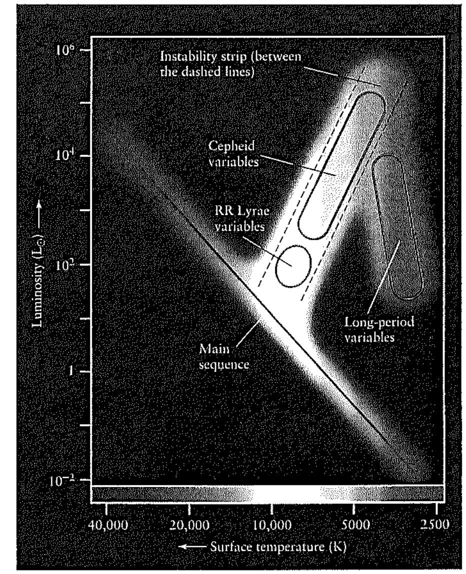
Astronomers do not fully understand why some cool red giants become long-period variables. It is difficult to calculate accurate stellar models to describe such huge stars with extended, tenuous atmospheres.




WEB LINK 19.9

**Figure 19-16** R I V U X G

Mira—A Long-Period Variable Star Mira, or o (omicron) Ceti, is a variable star whose luminosity varies with a 332-day period. At its dimmest, as in **(a)** (photographed in December 1961), Mira is less than 1% as bright as when it is at maximum, as in **(b)** (January 1965). These brightness variations occur because Mira pulsates. (Lowell Observatory)




INTERACTIVE EXERCISE 19.2

**Figure 19-17**

Variable Stars on the H-R Diagram Pulsating variable stars are found in the upper right of the H-R diagram. Long-period variables like Mira are cool red giant stars that pulsate slowly, changing their brightness in a semi-regular fashion over months or years. Cepheid variables and RR Lyrae variables are located in the instability strip, which lies between the main sequence and the red-giant region. A star passing through this strip along its evolutionary track becomes unstable and pulsates.

## Cepheid Variables

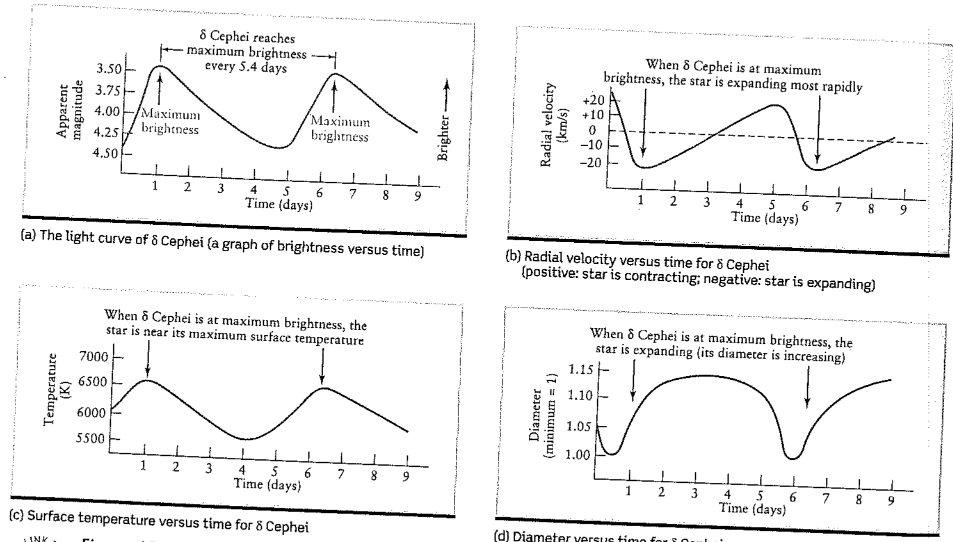
Astronomers have a much better understanding of other pulsating stars, called **Cepheid variables**, or simply Cepheids. A Cepheid variable is recognized by the characteristic way in which its light output varies—rapid brightening followed by gradual dimming. They are named for δ (delta) Cephei, an example of this type of star discovered in 1784 by John Goodricke, a deaf, mute, 19-year-old English amateur astronomer. He found that at its most brilliant, δ Cephei is 2.3 times as bright as at its dimmest. The cycle of brightness variations repeats every 5.4 days. (Sadly, Goodricke paid for his discoveries with his life; he caught pneumonia while making his nightly observations and died before his twenty-second birthday.) The surface temperatures and luminosities of the Cepheid variables place them in the upper middle of the H-R diagram (see Figure 19-17).

**By studying variable stars, astronomers gain insight into late stages of stellar evolution**

After core helium fusion begins, mature stars move across the middle of the H-R diagram. Figure 19-9*a* shows the evolutionary tracks of high-mass stars crisscrossing the H-R diagram. Post-helium-flash stars of moderate mass also cross the middle of the H-R diagram between the red-giant region and the horizontal branch.

During these transitions across the H-R diagram, a star can become unstable and pulsate. In fact, there is a region on the H-R diagram between the upper main sequence and the red-giant branch called the **instability strip** (see Figure 19-17). When an evolving star passes through this region, the star pulsates and its brightness varies periodically. Figure 19-18*a* shows the brightness variations of δ Cephei, which lies within the instability strip.

A Cepheid variable brightens and fades because the star's outer envelope cyclically expands and contracts. The first to observe this was the Russian astronomer Aristarkh Belopol'skii, who noticed in 1894 that spectral lines in the spectrum of δ Cephei shift back and forth with the same 5.4-day period as that of the magnitude variations. From the Doppler effect, we can




(a) The light curve of δ Cephei (a graph of brightness versus time)

(b) Radial velocity versus time for δ Cephei (positive: star is contracting; negative: star is expanding)

(c) Surface temperature versus time for δ Cephei

(d) Diameter versus time for δ Cephei

**Figure 19-18**

δ Cephei—A Pulsating Star **(a)** As δ Cephei pulsates, it brightens quickly (the light curve moves upward sharply) but fades more slowly (the curve declines more gently). The increases and decreases in brightness are nearly in step with variations in **(b)**, the star's radial velocity (positive when the star contracts and the surface moves away from us, negative when the star expands and the surface approaches us), as well as in **(c)**, the star's surface temperature. **(d)** The star is still expanding when it is at its brightest and hottest (compare with parts a and b).



translate these wavelength shifts into radial velocities and draw a velocity curve (Figure 19-18*b*). Negative velocities mean that the star's surface is expanding toward us; positive velocities mean that the star's surface is receding. Note that the light curve and velocity curve are mirror images of each other. The star is brighter than average while it is expanding and dimmer than average while contracting.

When a Cepheid variable pulsates, the star's surface oscillates up and down like a spring. During these cyclical expansions and contractions, the star's gases alternately heat up and cool down. Figure 19-18*c* shows the resulting changes in the star's surface temperature. Figure 19-18*d* graphs the periodic changes in the star's diameter.

Just as a bouncing ball eventually comes to rest, a pulsating star would soon stop pulsating without something to keep its oscillations going. In 1914, the British astronomer Arthur Eddington suggested that a Cepheid pulsates because the star is more opaque when compressed than when expanded. When the star is compressed, trapped heat increases the internal pressure, which pushes the star's surface outward. When the star expands, the heat escapes, the internal pressure drops, and the star's surface falls inward.

In the 1960s, the American astronomer John Cox followed up on Eddington's idea and proved that helium is what keeps Cepheids pulsating. Normally, when a star's helium is compressed, the gas increases in temperature and becomes more transparent. But in certain layers near the star's surface, compression may ionize helium (remove one of its electrons) instead of raising its temperature. Ionized helium gas is quite opaque, so these layers effectively trap heat and make the star expand, as Eddington suggested. This expansion cools the outer layers and makes the helium ions recombine with electrons, which makes the gas more transparent and releases the trapped energy. The star's surface then falls inward, recompressing the helium, and the cycle begins all over again.

**CAUTION!** In our discussion of the behavior of gases (see Box 19-1, Section 19-1, and Section 19-3) we saw that a gas cools when it expands and heats up when it is compressed. Hence, you would expect that the gases in a pulsating star like δ Cephei would reach their maximum temperature when the star is at its smallest diameter, so that the gases are most compressed. The hotter the gas, the more brightly it glows, so δ Cephei should also have its maximum brightness when its diameter is smallest. But Figure 19-18 shows that the star's brightness and the temperature of the gases at the surface reach their maximum values when the star is expanding, some time *after* the star has contracted to its smallest diameter. How can this be? The explanation is again related to how opaque the gases are inside the star. The rate at which energy is emitted from the central regions of the star is indeed greatest when the star is at its minimum diameter, but the opaque gases in the star's outer layers impede the flow of energy to the surface. Hence, δ Cephei reaches its maximum brightness and maximum surface temperature about half a day after the star is at its smallest size.

Cepheid variables are important because they have two properties that allow astronomers to determine the distances to very remote objects. First, Cepheids can be seen even at distances of millions of parsecs, because they are very luminous, ranging from a few hundred times solar luminosity to more than $10^4$ $L_\odot$. Second, there is a direct relationship between a Cepheid's period and its average luminosity: The dimmest Cepheid variables pulsate rapidly, with periods of 1 to 2 days, while the most luminous Cepheids pulsate with much slower periods of about 100 days.

Figure 19-19 shows this **period-luminosity relation.** By measuring the period of a distant Cepheid's brightness variations and using a graph like Figure 19-19, an astronomer can determine the star's luminosity. By also measuring the star's apparent brightness, the distance to the Cepheid can be found by using the inverse-square law (see Section 17-2). By applying the period-luminosity relation in this way to Cepheids in other galaxies, astronomers have been able to calculate the distances to those galaxies with great accuracy. (Box 17-2 gives an example of such a calculation.) As we will see in Chapters 24 and 26, such measurements play an important role in determining the overall size and structure of the universe.

How a Cepheid pulsates depends on the amount of heavy elements in the star's outer layers, because even trace amounts of these elements can have a large effect on how opaque the stellar gases are. Hence, Cepheids are classified according to their metal content. If the star is a metal-rich, Population I star, it is called a **Type I Cepheid;** if it is a metal-poor, Population II star, it is called

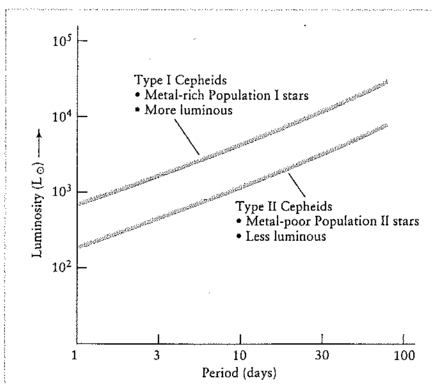



**Figure 19-19**

Period-Luminosity Relations for Cepheids The greater the average luminosity of a Cepheid variable, the longer its period and the slower its pulsations. Note that there are actually two distinct period-luminosity relations—one for Type I Cepheids and one for the less luminous Type II Cepheids. (Adapted from H. C. Arp) a **Type II Cepheid.** As Figure 19-19 shows, these two types of Cepheids exhibit different period-luminosity relations. In order to know which period-luminosity relation to apply to a given Cepheid, an astronomer must determine the star's metal content from its spectrum (see Figure 19-15).

The evolutionary tracks of mature, high-mass stars pass back and forth through the upper end of the instability strip on the H-R diagram. These stars become Cepheids when helium ionization occurs at just the right depth to drive the pulsations. For stars on the high-temperature (left) side of the instability strip, helium ionization occurs too close to the surface and involves only an insignificant fraction of the star's mass. For stars on the cool (right) side of the instability strip, convection in the star's outer layers prevents the storage of the energy needed to drive the pulsations. Thus, Cepheids exist only in a narrow temperature range on the H-R diagram.

### RR Lyrae Variables

Stars of lower mass do not become Cepheids. Instead, after leaving the main sequence, becoming red giants, and undergoing the helium flash, their evolutionary tracks pass through the lower end of the instability strip as they move along the horizontal branch. Some of these stars become **RR Lyrae variables,** named for their prototype in the constellation Lyra (the Harp). RR Lyrae variables all have periods shorter than one day and roughly the same average luminosity as horizontal-branch stars, about 100 $L_\odot$. In fact, the RR Lyrae region of the instability strip (see Figure 19-17) is actually a segment of the horizontal branch. RR Lyrae stars are all metal-poor, Population II stars. Many have been found in globular clusters, and they have been used to determine the distances to those clusters in the same way that Cepheids are used to find the distances to other galaxies. In Chapter 23 we will see how RR Lyrae stars helped astronomers determine the size of the Milky Way Galaxy.

In some cases the expansion speed of a pulsating star exceeds the star's escape speed. When this happens, the star's outer layers are ejected completely. We will see in Chapter 20 that dying stars eject significant amounts of mass in this way, renewing and enriching the interstellar medium for future generations of stars.

## 19-7 Mass transfer can affect the evolution of stars in a close binary system

We have outlined what happens when a main-sequence star evolves into a red giant. What we have ignored is that more than half of all stars are members of multiple-star systems, including binaries. If the stars in such a system are widely separated, the individual stars follow the same course of evolution as if they were isolated. In a **close binary,** however, when one star expands to become a red giant, its outer layers can be gravitationally captured by the nearby companion star. In other words, a bloated red giant in a close binary system can dump gas onto its companion, a process called **mass transfer.**

### Roche Lobes and Lagrangian Points

Our modern understanding of mass transfer in close binaries is based on the work of the French mathematician Edouard Roche. In the mid-1800s, Roche studied how rotation and mutual tidal interaction affect the stars in a binary system. Tidal forces cause the two stars in a close binary to keep the same sides facing each other, just as our Moon keeps its same side facing Earth (see Section 4-8). But because stars are gaseous, not solid, rotation and tidal forces can have significant effects on their shapes.

**If the stars in a binary system are sufficiently close, tidal forces can pull gases off one star and onto the other**

In widely separated binaries, the stars are so far apart that tidal effects are small, and, therefore, the stars are nearly perfect spheres. In close binaries, where the separation between the stars is not much greater than their sizes, tidal effects are strong, causing the stars to be somewhat egg-shaped.

Roche discovered a mathematical surface that marks the gravitational domain of each star in a close binary. (This surface is not a real physical one, like the surface of a balloon, but a mathematical construct.) Figure 19-20*a* shows the outline of this surface as a dashed line. The two halves of this surface, each of which encloses one of the stars, are known as **Roche lobes.** The more massive star is always located inside the larger Roche lobe. If gas from a star leaks over its Roche lobe, it is no longer bound by gravity to that star. This escaped gas is free either to fall onto the companion star or to escape from the binary system.

The point where the two Roche lobes touch, called the **inner Lagrangian point,** is a kind of balance point between the two stars in a binary. It is here that the effects of gravity and rotation cancel each other. When mass transfer occurs in a close binary, gases flow through the inner Lagrangian point from one star to the other.

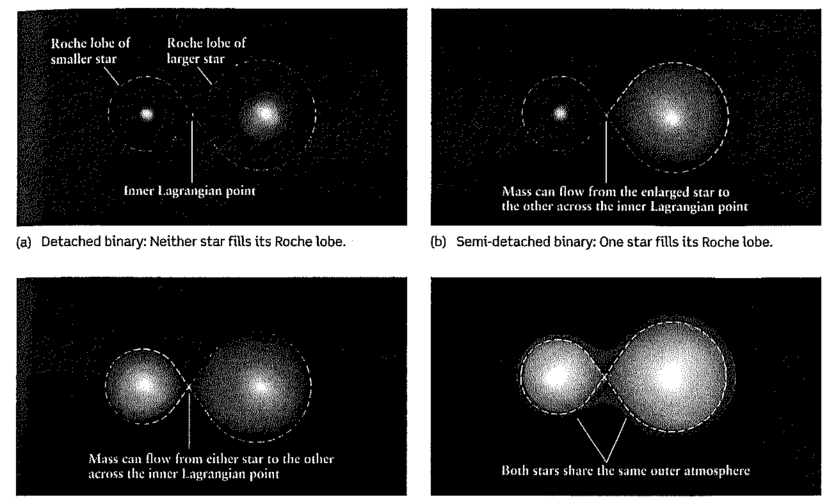
In many binaries, the stars are so far apart that even during their red-giant stages the surfaces of the stars remain well inside their Roche lobes. As a result, little mass transfer can occur and each star lives out its life as if it were single and isolated. A binary system of this kind is called a **detached binary** (Figure 19-20*a*).

However, if the two stars are close enough, when one star expands to become a red giant, it may fill or overflow its Roche lobe. Such a system is called a **semidetached binary** (Figure 19-20*b*). If both stars fill their Roche lobes, the two stars actually touch and the system is called a **contact binary** (Figure 19-20*c*). It is quite unlikely, however, that both stars exactly fill their Roche lobes at the same time. (This would only be the case if the two stars had identical masses, so that they both evolved at exactly the same rate.) It is more likely that they overflow their Roche lobes, giving rise to a common envelope of gas. Such a system is called an **overcontact binary** (Figure 19-20*d*).

### Observations of Mass Transfer

The binary star system Algol (from an Arabic term for "demon") provided the first clear evidence of mass transfer in close binaries. Also called β (beta) Persei, Algol can easily be seen with the naked






(a) Detached binary: Neither star fills its Roche lobe.

(b) Semi-detached binary: One star fills its Roche lobe.

(c) Contact binary: Both stars fill their Roche lobes.

(d) Overcontact binary: Both stars overfill their Roche lobes.

INTERACTIVE EXERCISE 19.3

**Figure 19-20**

Close Binary Star Systems The gravitational domain of a star in a close binary system is called its Roche lobe. The two Roche lobes meet at the inner Lagrangian point. The sizes of the stars relative to their Roche lobes determine whether the system is (a) a detached binary, (b) a semidetached binary, (c) a contact binary, or (d) an overcontact binary.

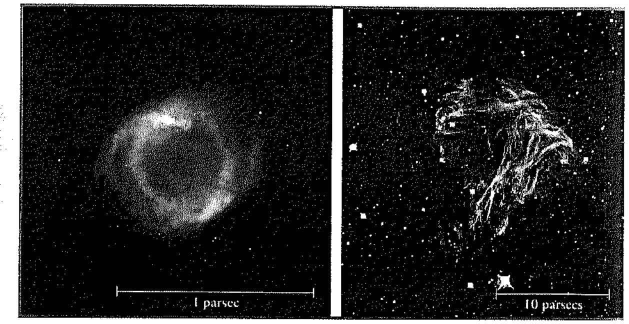
eye. Ancient astronomers knew that Algol varies periodically in brightness by a factor of more than 2. In 1782, John Goodricke (the discoverer of δ Cephei's variability) first suggested that these brightness variations take place because Algol is an *eclipsing* binary. (We discussed this type of binary in Section 17-11.) The orbital plane of the two stars that make up the binary system happens to be nearly edge-on to our line of sight, so one star periodically eclipses the other. Algol's light curve (Figure 19-21*a*) and spectrum show that Goodricke's brilliant hypothesis is correct, and that Algol is a semidetached binary. The detached star (on the right in Figure 19-21*a*) is a luminous blue main-sequence star, while its less massive companion is a dimmer red giant that fills its Roche lobe.

**CAUTION!** According to stellar evolution theory, the more massive a star, the more rapidly it should evolve. Since the two stars in a binary system form simultaneously and thus are the same age, the more massive star should become a red giant before the less massive one. But in Algol and similar binaries, the *more* massive star (on the right in Figure 19-21*a*) is still on the main sequence, whereas the *less* massive star (on the left in Figure 19-21*a*) has evolved to become a red giant. How can we explain this apparent contradiction? The answer is that the red giant in Algol-type binaries was *originally* the more massive star. As it left the main sequence to become a red giant, this star expanded until it overflowed its Roche lobe and dumped gas onto its originally less massive companion. Because of the resulting mass transfer, that companion (which is still on the main sequence) became the more massive star.

Mass transfer is also important in another class of semidetached binaries, called β (beta) Lyrae variables, after their prototype in the constellation Lyra. As with Algol, the less massive star in β Lyrae (on the left in Figure 19-21*b*) fills its Roche lobe. Unlike Algol, however, the more massive detached star (on the right in Figure 19-21*b*) is the dimmer of the two stars. Apparently, this detached star is enveloped in a rotating *accretion disk* of gas captured from its bloated companion. This disk partially blocks the light coming from the detached star, making it appear dimmer.

What is the fate of an Algol or β Lyrae system? If the detached star is massive enough, it will evolve rapidly, expanding to also fill its Roche lobe. The result will be an overcontact binary in which the two stars share the gases of their outer layers. Such binaries are sometimes called W Ursae Majoris stars, after the prototype of this class (Figure 19-21*c*).

# 20 Stellar Evolution: The Deaths of Stars




(a) A planetary nebula (b) A supernova remnant

WEB LINK 20.1 R I V U X G

Left: The planetary nebula NGC 7293. (the Helix Nebula) Right: The supernova remnant LMC N49. (NASA, NOAO, ESA, the Hubble Helix Nebula Team, M. Meixner/STScI, and T. A. Rector/NRAO; NASA and the Hubble Heritage Team, STScI/AURA)

When a star of 0.4 solar mass or more reaches the end of its main-sequence lifetime and becomes a red giant, it has a compressed core and a bloated atmosphere. Finally, the star devours its remaining nuclear fuel and begins to die. As we'll learn in this chapter, the character of the star's death depends crucially on the value of its mass.

A star of relatively low mass—such as our own Sun—ends its evolution by gently expelling its outer layers into space. These ejected gases form a glowing cloud called a *planetary nebula* such as the one shown here in the left-hand image. The burned-out core that remains is called a *white dwarf*.

In contrast, a high-mass star ends its life in almost inconceivable violence. At the end of its short life, the core of such a star collapses suddenly, which triggers a powerful *supernova* explosion that can be as luminous as an entire galaxy of stars. A white dwarf, too, can become a supernova if it accretes gas from a companion star in a close binary system.

Thermonuclear reactions in supernovae produce a wide variety of heavy elements, which are ejected into the interstellar medium. (The supernova remnant shown here in the right-hand image is rich in these elements.) Such heavy elements are essential building blocks for terrestrial worlds like our Earth. Thus, the deaths of massive stars can provide the seeds for planets orbiting succeeding generations of stars.

## 20-1 Stars of between 0.4 and 4 solar masses go through two distinct red-giant stages

TUTORIAL 20.1

All main-sequence stars convert hydrogen to helium in their cores in a series of energy-releasing thermonuclear reactions. As we saw in Section 19-1, convection within a low-mass main-sequence star—a so-called *red dwarf* with a mass between 0.08 and 0.4 $M_\odot$—will eventually bring all of the star's hydrogen into the core. Over hundreds of billions of years a red dwarf evolves into an inert ball of helium. Convection is less important in main-sequence stars with masses greater than 0.4 $M_\odot$, so these stars are able to consume only the hydrogen that is present within the core. These stars of greater mass then leave the main sequence. Let's examine what happens next for a star of moderately low mass, between 0.4 and 4 $M_\odot$. One example for such a star is our own Sun, with a mass of 1 $M_\odot$. We'll begin by

## Learning Goals

*By reading the sections of this chapter, you will learn*

20-1 What kinds of thermonuclear reactions occur inside a star of moderately low mass as it ages

20-2 How evolving stars disperse carbon into the interstellar medium

20-3 How stars of moderately low mass eventually die

20-4 The nature of white dwarfs and how they are formed

20-5 What kinds of reactions occur inside a high-mass star as it ages

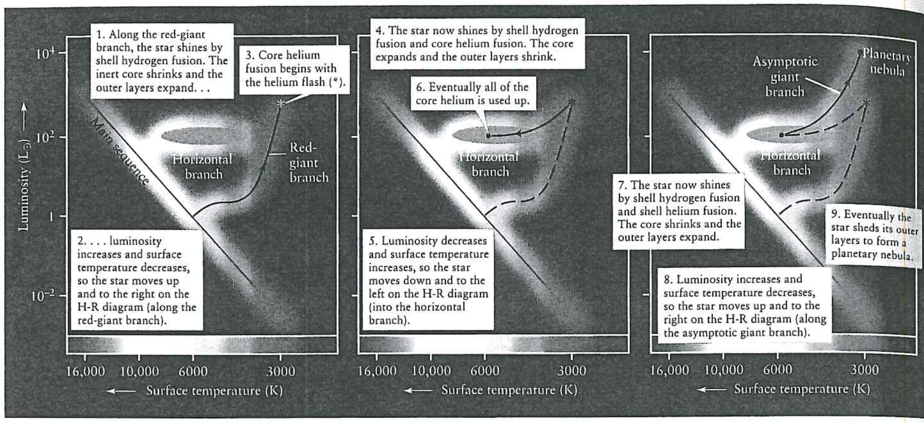
20-6 How high-mass stars explode and die

20-7 Why supernova SN 1987A was both important and unusual

20-8 What role neutrinos play in the death of a massive star

20-9 How white dwarfs in close binary systems can explode

20-10 What remains after a supernova explosion




(a) Before the helium flash: A red-giant star (b) After the helium flash: A horizontal-branch star (c) After core helium fusion ends: An AGB star

## Figure 20-1

The Post-Main-Sequence Evolution of a 1-$M_\odot$ Star These H-R diagrams show the evolutionary track of a star like the Sun as it goes through the stages of being **(a)** a red-giant star, **(b)** a horizontal-branch star, and **(c)** an asymptotic giant branch (AGB) star. The star eventually evolves into a planetary nebula (described in Section 20-3).

reviewing what we learned in Chapter 19 about the first stages of post–main-sequence evolution for such a star. (Later in this chapter we'll study the evolution of more massive stars.)

### The Red-Giant and Horizontal-Branch Stages: A Review

We can describe a star's post–main-sequence evolution using an evolutionary track on a H-R diagram. Figure 20-1 shows the track for a 1-$M_\odot$ star like the Sun. Once core hydrogen fusion ceases, the core shrinks, heating the surrounding hydrogen and triggering shell hydrogen fusion. The new outpouring of energy causes the star's outer layers to expand and cool, and the star becomes a red giant. As the luminosity increases and the surface temperature drops, the post–main-sequence star moves up and to the right along the **red-giant branch** on an H-R diagram (Figure 20-1*a*).

Next, the helium-rich core of the star shrinks and heats until eventually **core helium fusion** begins. This second post–main-sequence stage begins gradually in stars more massive than about 2–3 $M_\odot$, but for less massive stars it comes suddenly—in a *helium flash*. During core helium fusion, the surrounding hydrogen-fusing shell still provides most of the red giant's luminosity.

As we learned in Section 19-3, the core expands when core helium fusion begins, which makes the core cool down a bit. (We saw in Box 19-1 that letting a gas expand tends to lower its temperature, while compressing a gas tends to increase its temperature.) The cooling of the core also cools the surrounding hydrogen-fusing shell, so that the shell releases energy more slowly. Hence, the luminosity goes down a bit after core helium fusion begins.

The slower rate of energy release also lets the star's outer layers contract. As they contract, they heat up, so the star's surface temperature increases and its evolutionary track moves to the left on the H-R diagram in Figure 20-1*b*. The luminosity changes relatively little during this stage, so the evolutionary track moves almost horizontally, along a path called the **horizontal branch**. Horizontal-branch stars have helium-fusing cores surrounded by hydrogen-fusing shells. Figure 19-12 shows horizontal-branch stars in a globular cluster, and Figure 19-8 shows the evolution of the luminosity of a 1-$M_\odot$ star up to this point in its history.

### AGB Stars: The Second Red-Giant Stage

Helium fusion produces nuclei of carbon and oxygen. After about a hundred million ($10^8$) years of core helium fusion, essentially all the helium in the core of a 1-$M_\odot$ star has been converted into carbon and oxygen, and the fusion of helium in the core ceases. (This corresponds to the right-hand end of the graph in Figure 19-8.) Without thermonuclear reactions to maintain the core's internal pressure, the core again contracts, until it is stopped by degenerate-electron pressure (described in Section 19-3). This contraction releases heat into the surrounding helium-rich gases, and a new stage of helium fusion begins in a thin shell around the core. This process is called **shell helium fusion.**

**Stars like the Sun go through a second red-giant phase, during which helium fusion takes place in a shell around an inert core**

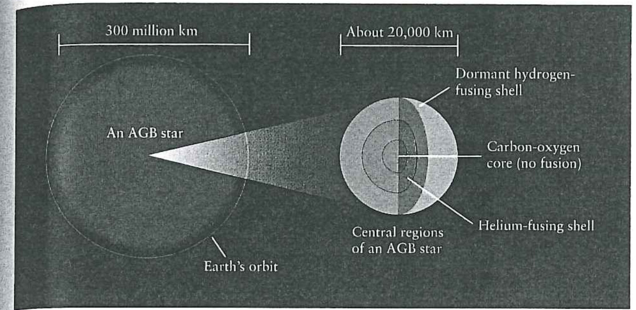
History now repeats itself—the star enters a *second* red-giant phase. A star first becomes a red giant at the end of its main-sequence lifetime, when the outpouring of energy from shell hydrogen fusion makes the star's outer layers expand and cool. In the same way, the outpouring of energy from shell *helium* fusion causes the outer layers to expand again. The low-mass star ascends into the red-giant region of the H-R diagram for a second time (Figure 20-1*c*), but now with even greater luminosity than during its first red-giant phase.

Stars in this second red-giant phase are commonly called **asymptotic giant branch stars,** or **AGB stars,** and their evolutionary tracks follow what is called the **asymptotic giant branch.** (*Asymptotic* means "approaching"; the name means that a star on the asymptotic giant branch approaches the red-giant branch from the left on an H-R diagram.)

When a low-mass star first becomes an AGB star, it consists of an inert, degenerate carbon-oxygen core and a helium-fusing shell, both inside a hydrogen-fusing shell, all within a volume not much larger than Earth. This small, dense central region is surrounded by an enormous hydrogen-rich envelope about as big as Earth's orbit around the Sun. After a while, the expansion of the star's outer layers causes the hydrogen-fusing shell to also expand and cool, and thermonuclear reactions in this shell temporarily cease. This leaves the aging star's structure as shown in Figure 20-2.

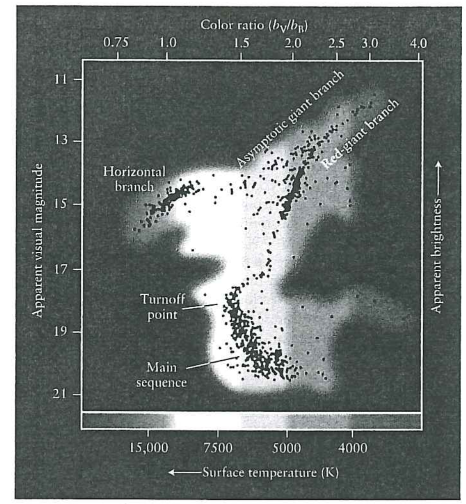
We saw in Section 19-1 that the more massive a star, the shorter the amount of time it remains on the main sequence. Similarly, the greater the star's mass, the more rapidly it goes through the stages of post–main-sequence evolution. Hence, we can see all of these stages by studying star clusters, which contain stars that are all the same age but that have a range of masses (see Section 19-4). Figure 20-3 shows a color-magnitude diagram for the globular cluster M55, which is at least 13 billion years old. The least massive stars in this cluster are still on the main sequence. Progressively more massive stars have evolved to the red giant branch, the horizontal branch, and the asymptotic giant branch.






## Figure 20-2

The Structure of an Old, Moderately Low-Mass AGB Star Near the end of its life, a star like the Sun becomes an immense, red, asymptotic giant branch (AGB) star. The star's inert core, active helium-fusing shell, and dormant hydrogen-fusing shell are all contained within a volume roughly the size of Earth. Thermonuclear reactions in the helium-fusing shell are so rapid that the star's luminosity is thousands of times that of the present-day Sun. (The relative sizes of the shells in the star's interior are not shown to scale.)




## Figure 20-3

Stellar Evolution in a Globular Cluster In the old globular cluster M55, stars with masses less than about 0.8 $M_\odot$ are still on the main sequence, converting hydrogen into helium in their cores. Slightly more massive stars have consumed their core hydrogen and are ascending the red-giant branch; even more massive stars have begun helium core fusion and are found on the horizontal branch. The most massive stars (which still have less than 4 $M_\odot$) have consumed all the helium in their cores and are ascending the asymptotic giant branch. (Compare with Figure 21-11.) (Adapted from D. Schade, D. VandenBerg, and F. Hartwick)

A 1-$M_\odot$ AGB star can reach a maximum luminosity of nearly $10^4$ $L_\odot$, as compared with approximately $10^3$ $L_\odot$ when it reached the helium flash and a relatively paltry 1 $L_\odot$ during its main-sequence lifetime. When the Sun becomes an AGB star some 7.8 billion years from now, this tremendous increase in luminosity will cause Mars and the Jovian planets to largely evaporate away. The Sun's bloated outer layers will reach to Earth's orbit. Mercury and perhaps Venus will simply be swallowed whole.

## 20-2 Dredge-ups bring the products of nuclear fusion to a giant star's surface

As we saw in Section 16-2, energy is transported outward from a star's core by one of two processes—radiative diffusion or convection. The first is the passage of energy in the form of electromagnetic radiation, and it dominates only when a star's gases are relatively transparent. The second involves up-and-down movement of the star's gases. Convection plays a very important role in giant stars, and it helps supply the cosmos with the elements essential to life.

### Convection, Dredge-ups, and Carbon Stars

ANIMATION 20.2

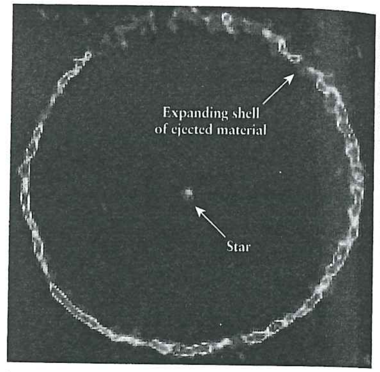
In the Sun, convection dominates only the outer layers, from around 0.71 solar radius (measured from the center of the Sun) up to the photosphere (recall Figure 16-4). During the final stages of a star's life, however, the convective zone can become so broad that it extends down to the star's core. At these times, convection can "dredge up" the heavy elements produced in and around the core by thermonuclear fusion, transporting them all the way to the star's surface.

LOOKING DEEPER 20.1

The *first* **dredge-up** takes place after core hydrogen fusion stops, when the star becomes a red giant for the first time. Convection dips so deeply into the star that material processed by the CNO cycle of hydrogen fusion (see Section 16-1) is carried up to the star's surface, changing the relative abundances of carbon, nitrogen, and oxygen. A *second* dredge-up occurs after core helium fusion ceases, further altering the abundances of carbon, nitrogen, and oxygen. Still later, during the AGB stage, a *third* dredge-up can occur if the star has a mass greater than about 2 $M_\odot$. This third dredge-up transports large amounts of freshly synthesized carbon to the star's surface, and the star's spectrum thus exhibits prominent absorption bands of carbon-rich molecules like $C_2$, CH, and CN. For this reason, an AGB star that has undergone a third dredge-up is called a **carbon star**.

All AGB stars have very strong stellar winds that cause them to lose mass at very high rates, up to $10^{-4}$ $M_\odot$ per year (a thousand times greater than that of a red giant, and $10^{10}$ times greater than the rate at which our present-day Sun loses mass). The surface temperature of AGB stars is relatively low, around 3000 K, so any ejected carbon-rich molecules can condense to form tiny grains of soot. Indeed, carbon stars are commonly found to be obscured in sooty cocoons of ejected matter (Figure 20-4).

**The carbon that forms the basis of all life on Earth was ejected billions of years ago from giant stars**




WEB LINK 20.3

**Figure 20-4** R I V U X G
A Carbon Star TT Cygni is an AGB star in the constellation Cygnus that ejects some of its carbon-rich outer layers into space. Some of the ejected carbon combines with oxygen to form molecules of carbon monoxide (CO), whose emissions can be detected with a radio telescope. This radio image shows the CO emissions from a shell of material that TT Cygni ejected some 7000 years ago. Over that time, the shell has expanded to a diameter of about ½ light-year. (H. Olofsson, Stockholm Observatory, et al./NASA)

Carbon stars are important because they enrich the interstellar medium with carbon and some nitrogen and oxygen. The triple alpha process that occurs in helium fusion is the *only* way that carbon can be made, and carbon stars are the primary avenue by which this element is dispersed into interstellar space. Indeed, most of the carbon in your body was produced long ago inside a star by the triple alpha process (see Section 19-3). This carbon was later dredged up to the star's surface and ejected into space. Some 4.56 billion years ago a clump of the interstellar medium which contained this carbon coalesced into the solar nebula from which our Earth—and all of the life on it—eventually formed. In this sense you can think of your body as containing "recycled" material—substances that were once in the heart of a star that formed and evolved long before our solar system existed.

## 20-3 Stars of moderately low mass die by gently ejecting their outer layers, creating planetary nebulae

For a star that began with a moderately low mass (between about 0.4 and 4 $M_\odot$), the AGB stage in its evolution is a dramatic turning point. Before this stage, a star loses mass only gradually through steady stellar winds. But as it evolves during its AGB stage, a star divests itself completely of its outer layers. The aging star undergoes a series of bursts in luminosity, and in each burst it ejects a shell of material into space. (The shell around the AGB star TT Cygni, shown in Figure 20-4, was probably created in this way.) Eventually, all that remains of a low-mass star is a fiercely hot, exposed core, surrounded by glowing shells of ejected gas. This late stage in the life of a star is called a **planetary nebula**. The left-hand image on the opening page of this chapter shows one such planetary nebula, called the Ring Nebula for its shape.

**An aging AGB star casts off much of the mass that it possesses and makes a beautiful glowing nebula.**



### Making a Planetary Nebula

To understand how an AGB star can eject its outer layers in shells, consider the internal structure of such a star as shown in Figure 20-2. As the helium in the helium-fusing shell is used up, the pressure that holds up the dormant hydrogen-fusing shell decreases. Hence, the dormant hydrogen shell contracts and heats up, and hydrogen fusion begins anew. This revitalized hydrogen fusion creates helium, which rains downward onto the temporarily dormant helium-fusing shell. As the helium shell gains mass, it shrinks and heats up. When the temperature of the helium shell reaches a certain critical value, it reignites in a **helium shell flash** that is similar to (but less intense than) the helium flash that occurred earlier in the evolution of a low-mass star (see Section 19-3). The released energy pushes the hydrogen-fusing shell outward, making it cool off, so that hydrogen fusion ceases and this shell again becomes dormant. The process then starts over again.

When a helium shell flash occurs, the luminosity of an AGB star increases substantially in a relatively short-lived burst called a **thermal pulse.** Figure 20-5, which is based on a theoretical calculation of the evolution of a 1-$M_\odot$ star, shows that thermal pulses begin when the star is about 12.365 billion years old. The calculations predict that thermal pulses occur at ever-shorter intervals of about 100,000 years.

During these thermal pulses, the dying star's outer layers can separate completely from its carbon-oxygen core. As the ejected material expands into space, dust grains condense out of the cooling gases. Radiation pressure from the star's hot, burned-out core acts on the specks of dust, propelling them further outward, and the star sheds its outer layers altogether. In this way an aging 1-$M_\odot$ star loses as much as 40% of its mass. More massive stars eject even greater fractions of their original mass.

As a dying star ejects its outer layers, the star's hot core becomes exposed. With a surface temperature of about 100,000 K, this exposed core emits ultraviolet radiation intense enough to ionize and excite the expanding shell of ejected gases. These gases therefore glow and emit visible light through the process of fluorescence (see Box 18-1), producing a planetary nebula like those shown in Figure 20-6.

**CAUTION!** Despite their name, planetary nebulae have nothing to do with planets. This misleading term was introduced in the nineteenth century because these glowing objects looked like distant Jovian planets—a small colored blur—when viewed through the small telescopes then available. The difference between planets and planetary nebulae first became obvious with the advent of spectroscopy: Planets have *absorption* line spectra (see Section 7-3), but the excited gases of planetary nebulae have *emission* line spectra.

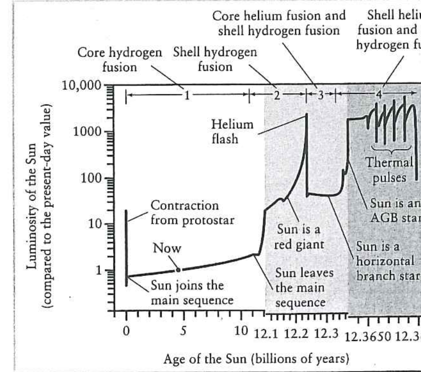



**Figure 20-5**
Further Stages in the Evolution of the Sun This diagram, which shows how the luminosity of the Sun (a 1-$M_\odot$ star) changes over time, is an extension of Figure 19-8. We use different scales for the final stages because the evolution is so rapid. During the AGB stage there are brief periods of runaway helium fusion, causing spikes in luminosity called thermal pulses. (Adapted from Mark A. Garlick, based on calculations by I.-Juliana Sackmann and Kathleen E. Kramer)

### The Properties of Planetary Nebulae

Planetary nebulae are quite common. Astronomers estimate that there are 20,000 to 50,000 planetary nebulae in our Galaxy alone. Many planetary nebulae, such as those in Figure 20-6, are more or less spherical in shape. This shape is a result of the symmetrical way in which the gases were ejected. But if the rate of expansion is not the same in all directions, the resulting nebula can take on an hourglass or dumbbell appearance (Figure 20-7).

Spectroscopic observations of planetary nebulae show emission lines of ionized hydrogen, oxygen, and nitrogen. From the Doppler shifts of these lines, astronomers have concluded that the expanding shell of gas moves outward from a dying star at speeds from 10 to 30 km/s. For a shell expanding at such speeds to have attained the typical diameter of a planetary nebula, about 1 light-year, it must have begun expanding about 10,000 years ago. Thus, by astronomical standards, the planetary nebulae we see today were created only very recently.

We do not observe planetary nebulae that are more than about 50,000 years old. After this length of time, the shell has spread out so far from the cooling central star that its gases cease to glow and simply fade from view. The nebula's gases then mix with the surrounding interstellar medium.

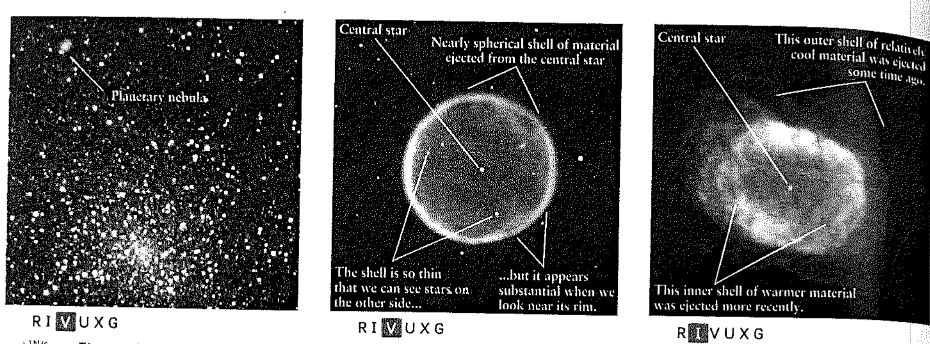



**Figure 20-6**

Planetary Nebulae (a) The pinkish blob is a planetary nebula surrounding a star in the globular cluster M15, about 10,000 pc (33,000 ly) from Earth in the constellation Pegasus. (b) The planetary nebula Abell 39 lies about 2200 pc (7000 ly) from Earth in the constellation Hercules. The almost perfectly spherical shell that comprises the nebula is about 1.5 pc (5 ly) in diameter; the thickness of the shell is only about 0.1 pc (0.3 ly). (c) This infrared image of the planetary nebula NGC 7027 suggests a more complex evolutionary history than that of Abell 39. NGC 7027 is about 900 pc (3000 ly) from Earth in the constellation Cygnus and is roughly 14,000 AU across. (a: NASA/Hubble Heritage Team, STScI/AURA; b: WIYN/NOAO/NSF; c: William B. Latter, SIRTF Science Center/Caltech, and NASA)

Astronomers estimate that all the planetary nebulae in the Galaxy return a total of about 5 $M_{\odot}$ to the interstellar medium each year. This amount is about 15% of all the matter expelled by all the various sorts of stars in the Galaxy each year. Because this contribution is so significant, and because the ejected material includes heavier elements (metals) manufactured within a nebula's central star, planetary nebulae play an important role in the chemical evolution of the Galaxy as a whole.

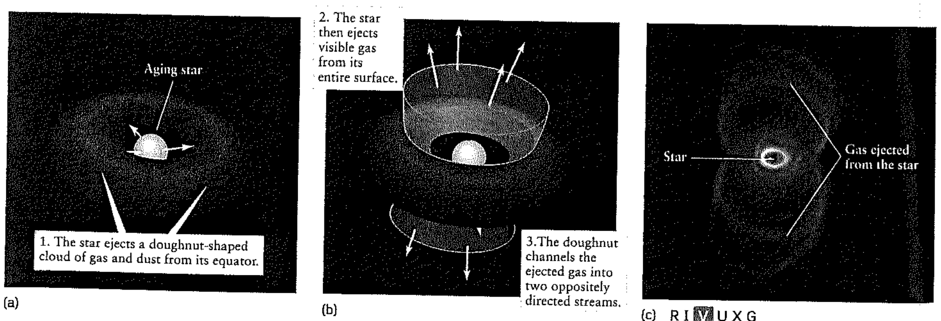



**Figure 20-7**

Making an Elongated Planetary Nebula (a), (b) These illustrations show one proposed explanation for why many planetary nebulae have an elongated shape. (c) The planetary nebula MyCn18, shown here in false color, may have acquired its elongated shape in this way. It lies some 2500 pc (8000 ly) from Earth in the constellation Musca (the Fly). (R. Sahai and J. Trauger, Jet Propulsion Laboratory; the WFPC-2 Science Team; and NASA)

## 20-4 The burned-out core of a moderately low-mass star cools and contracts until it becomes a white dwarf

We have seen that after a moderately low-mass star (from about 0.4 to about 4 solar masses) consumes all the hydrogen in its core, it is able to ignite thermonuclear reactions that convert helium to carbon and oxygen. Given sufficiently high temperature and pressure, carbon and oxygen can also undergo fusion reactions that release energy. But for such a moderately low-mass star, the core temperature and pressure never reach the extremely high values needed for these reactions to take place. Instead, as we have seen, the process of mass ejection just strips away the star's outer layers and leaves behind the hot carbon-oxygen core. With no thermonuclear reactions taking place, the core simply cools down like a dying ember. Such a burnt-out relic of a star's former glory is called a **white dwarf.** Such white dwarfs prove to have exotic physical properties that are wholly unlike any object found on Earth.

**CAUTION!** Unfortunately, the word *dwarf* is used in astronomy for several very different kinds of small objects. Here's a review of the three kinds that we have encountered so far in this book. A *white* dwarf is the relic that remains at the very end of the evolution of a star of initial mass between about 0.4 $M_{\odot}$ and 4 $M_{\odot}$. Thermonuclear reactions are no longer taking place in its interior; it emits light simply because it is still hot. A *red* dwarf, discussed in Section 19-1, is a cool main-sequence star with a mass between about 0.08 $M_{\odot}$ and 0.4 $M_{\odot}$. The energy emitted by a red dwarf in the form of light comes from its core, where fusion reactions convert hydrogen into helium. Finally, a *brown* dwarf (see Section 8-6 and Section 17-5) is an object like a main-sequence star but with a mass less than about 0.08 $M_{\odot}$. Because its mass is so small, its internal pressure and temperature are too low to sustain thermonuclear reactions. Instead, a brown dwarf emits light because it is slowly contracting, a process that releases energy (see Section 16-1). White dwarfs are comparable in size to Earth (see Section 17-7); by contrast, brown dwarfs are larger than the planet Jupiter, and red dwarfs are even larger.

### Properties of White Dwarfs

You might think that without thermonuclear reactions to provide internal heat and pressure, a white dwarf should keep on shrinking under the influence of its own gravity as it cools. Actually, however, a cooling white dwarf maintains its size, because the burnt-out stellar core is so dense that most of its electrons are degenerate (see Section 17-3). Thus, degenerate-electron pressure supports the star against further collapse. This pressure does not depend on temperature, so it continues to hold up the star even as the white dwarf cools and its temperature drops.

**A white dwarf is kept from collapsing by the pressure of its degenerate electrons**

Many white dwarfs are found in the solar neighborhood, but all are too faint to be seen with the naked eye. One of the first white dwarfs to be discovered is a companion to Sirius, the brightest star in the night sky. In 1844 the German astronomer Friedrich Bessel noticed that Sirius was moving back and forth slightly, as if it was being orbited by an unseen object. This companion, designated Sirius B (Figure 20-8), was first glimpsed in 1862 by the American astronomer Alvan Clark. Recent Hubble Space Telescope observations at ultraviolet wavelengths, where hot white dwarfs emit most of their light, show that the surface temperature of Sirius B is 25,200 K. (By contrast, the main-sequence star Sirius A has a surface temperature of 10,500 K, while the Sun's surface temperature is a relatively frosty 5800 K.)

Observations of white dwarfs in binary systems like Sirius allow astronomers to determine the mass, radius, and density of these stars (see Sections 17-9, 17-10, and 17-11). Such observations show that the density of the degenerate matter in a white dwarf is typically $10^9$ kg/m$^3$ (a million times denser than water). A teaspoonful of white dwarf matter brought to Earth would weigh nearly 5.5 tons—as much as an elephant!

### The Mass-Radius Relation for White Dwarfs

As we learned in Section 17-3, degenerate matter has a very different relationship between its pressure, density, and temperature

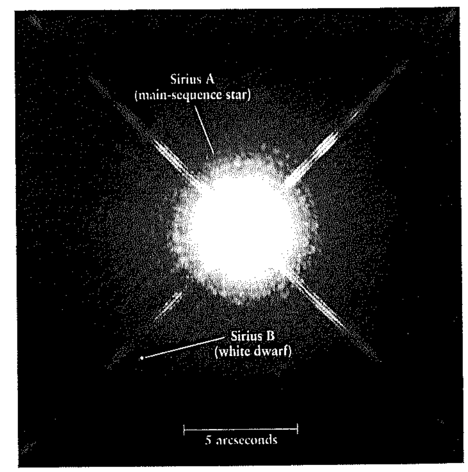



**Figure 20-8** R I V U X G

Sirius A and Its White Dwarf Companion Sirius, the brightest-appearing star in the sky, is actually a binary star: The secondary star, called Sirius B, is a white dwarf. In this Hubble Space Telescope image, Sirius B is almost obscured by the glare of the overexposed primary star, Sirius A, which is about $10^4$ times more luminous than Sirius B. The halo and rays around Sirius A are the result of optical effects within the telescope. (NASA; H. E. Bond and E. Nelan, STScI; M. Barstow and M. Burleigh, U. of Leicester; and J. B. Holberg, U. of Arizona)