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محاضرات علمية لمادة فيزياء الفلك / Astrophysics

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مدرس المادة

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Astrophysics Definition

Astrophysics can be defined as the study of the nature and physics of stars and star systems. It provides the theoretical framework for understanding astronomical observations.

At times astrophysics can be used to predict phenomena before they have even been observed by astronomers, such as **black holes**. The laboratory of outer space makes it possible to investigate large-scale physical processes that cannot be duplicated in a terrestrial laboratory.

Today astrophysics has within its reach the ability to bring about one of the greatest scientific achievements ever a unified understanding of the total evolutionary scheme of the universe. This remarkable revolution in astrophysics is happening now as a result of the confluence of two streams of technical development:

1- **Remote sensing:** Through the science of remote sensing we have acquired sensitive instruments capable of detecting and analyzing radiation across the whole range of the electromagnetic (EM) spectrum.

2- Spaceflight: Which lets astrophysicists place sophisticated remote sensing instruments above Earth's atmosphere. The wavelengths transmitted through the interstellar medium and arriving in the vicinity of near-Earth space are spread over approximately **24 decades** of the spectrum. However, the majority of this interesting EM radiation never reaches the surface of Earth because the terrestrial atmosphere effectively blocks such radiation across most of the spectrum. It should be remembered that **the visible and infrared** atmospheric windows occupy a spectral slice whose width is roughly **1 decade.**

Ground-based radio observatories can detect stellar radiation over a spectral range that adds about **5 more decades** to the range of observable frequencies, but the remaining **18 decades** of the spectrum are still blocked and are effectively invisible to astrophysicists on Earth's surface. Consequently, information that can be gathered by observers at the bottom of Earth's atmosphere represents only a small fraction of the total amount of information available concerning extraterrestrial objects. Sophisticated remote sensing instruments placed above Earth's atmosphere are now capable of sensing electromagnetic radiation over nearly the entire spectrum, and these instruments are rapidly changing our picture of the cosmos. See table and figure below.

For example, we previously thought that the interstellar medium was a fairly uniform collection of gas and dust, but space-borne ultraviolet telescopes have shown us that its structure is very inhomogenous and complex.

Table 1. The electromagnetic spectrum.

Region	Wavelength	Frequency (Hz)
Radio	> 1 mm	$< 3 \times 10^{11}$
Infrared	700 nm–1 mm	3×10^{11}–4.3×10^{14}
Visible	400–700 nm	4.3×10^{14}–7.5×10^{14}
Ultraviolet	10–400 nm	7.5×10^{14}–3×10^{16}
X-ray	0.1–10 nm	3×10^{16}–3×10^{18}
Gamma-ray	< 0.1 nm	$> 3 \times 10^{18}$

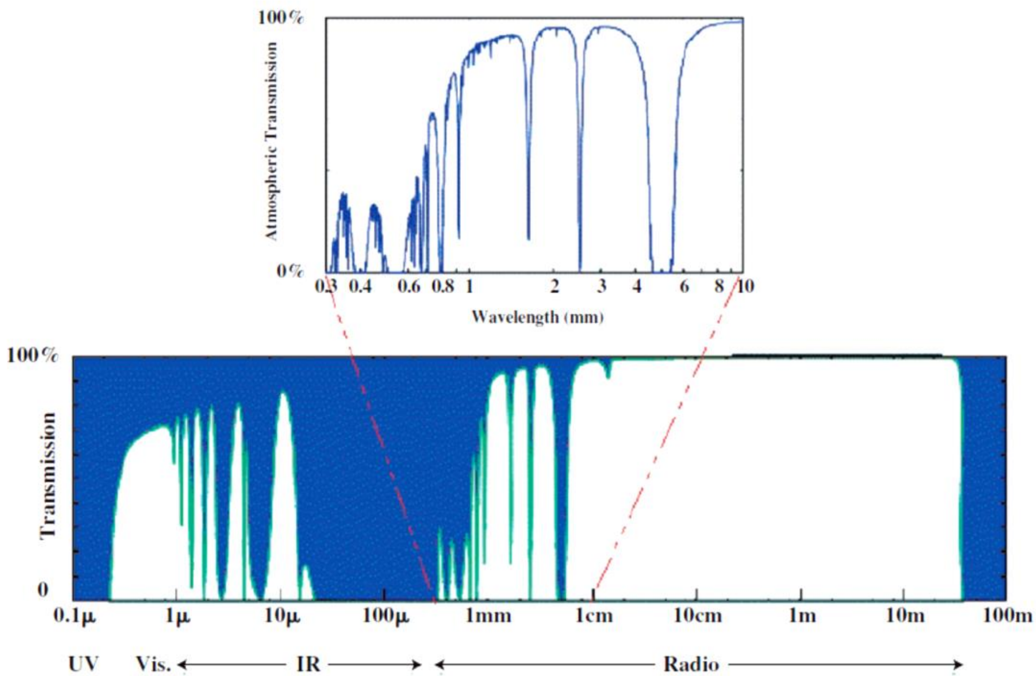


Figure 1: *Atmospheric transmission as a function of wavelength. The curve shows the fraction of transmission at each wavelength. Note the good transmission in the radio and visible parts of the spectrum. Also note a few narrow ranges, or "windows" of relatively good transmission in the infrared.*

The Tools of Astrophysics

Virtually all the information we receive about celestial objects comes to us through **observation of electromagnetic radiation**. **Cosmic ray particles** are an obvious and important exception, as are **extraterrestrial material samples** that have been returned to Earth (for example, lunar rocks). Each portion of the electromagnetic spectrum carries unique information about the physical conditions and processes in the universe. **Infrared radiation** reveals the presence of thermal emission from relatively cool objects; **ultraviolet and extreme ultraviolet radiation** may indicate thermal emission from very hot objects. Various types of violent events can lead to the production of **X-rays and gamma rays**. Although EM radiation varies over many decades of energy and wavelength, the basic principles of measurement are quite common to all regions of the spectrum.

The fundamental techniques used in astrophysics can be classified as imaging, spectrometry, photometry and polarimetry.

1- **Imaging:** provides basic information about the distribution of material in a celestial object, its overall structure, and, in some cases, its physical nature.

2- **Spectrometry:** is a measure of radiation intensity as a function of wavelength. It provides information on nuclear, atomic, and molecular phenomena occurring in and around the extraterrestrial object under observation.

3- **Photometry:** involves measuring radiation intensity as a function of time. It provides information about the time variations of physical processes within and around celestial objects, as well as their absolute intensities.

4- **polarimetry:** is a measurement of radiation intensity as a function of polarization angle. It provides information on ionized particles rotating in strong magnetic fields.

High-Energy Astrophysics

High-energy astrophysics encompasses the study of extraterrestrial X-rays, gamma rays, and energetic cosmic ray particles. Prior to space-based high-energy astrophysics, scientists believed that violent processes involving high energy emissions were rare in stellar and galactic evolution. Now, because of studies of extraterrestrial X-rays and gamma rays, we know that such processes are quite common rather than exceptional. **The observation of X-ray emissions has been very valuable in the study of high-energy events**, such as **mass transfer in binary star systems**, **interaction of supernova remnants with interstellar gas, and quasars** (whose energy source is presently unknown but believed to involve matter accreting to [falling into] a black hole). It is thought that **gamma rays** might be the missing link in understanding the physics of interesting high-energy objects such as **pulsars and black holes**. The study of **cosmic ray particles** provides important information about the physics of **nucleosynthesis** and about the **interactions of particles and strong magnetic fields**. **High-energy phenomena** that are suspected sources of cosmic rays include **supernovas, pulsars, radio galaxies, and quasars.**

X-Ray Astronomy

X-ray astronomy is the most advanced of the three high-energy astrophysics disciplines. Space-based X-ray observatories, such as NASA’s Chandra X-Ray Observatory, increase our understanding in the following areas:

(1) Stellar structure and evolution, including binary star systems, supernova remnants, pulsar and plasma effects, and relativity effects in intense gravitational fields.

(2) Large-scale galactic phenomena, including interstellar media and X-ray mapping of local galaxies.

(3) The nature of active galaxies, including spectral characteristics and the time variation of X-ray emissions from the nuclear or central regions of such galaxies.

(4) Rich clusters of galaxies, including X-ray background radiation and cosmology modeling.

Gamma-Ray Astronomy

Gamma rays consist of extremely energetic photons (that is, energies greater than 10^5 electron volts [eV]) and result from physical processes different than those associated with X-rays. The processes associated with gamma ray emissions in astrophysics include:

(1) The decay of radioactive nuclei.

(2) Cosmic ray interactions.

(3) Curvature radiation in extremely strong magnetic fields

(4) matter-antimatter annihilation.

Gamma-ray astronomy reveals the explosive, high-energy processes associated with such celestial phenomena as supernovas, exploding galaxies and quasars, pulsars, and black holes. Gamma-ray astronomy is especially significant because the gamma rays being observed can travel across our entire galaxy and even across most of the universe without suffering appreciable alteration or absorption. Therefore, these energetic gamma rays reach our solar system with the same characteristics, including directional and temporal features, as they started with at their sources, possibly many light-years distant and deep within regions or celestial objects opaque to other wavelengths. Consequently, gamma-ray astronomy provides information on extraterrestrial phenomena not observable at any other wavelength in the electromagnetic spectrum and on spectacularly energetic events that may have occurred far back in the evolutionary history of the universe.

Cosmic-Ray Astronomy

Cosmic rays are extremely energetic particles that extend in energy from one million (10^6) electron volts (eV) to more than 10^{20} eV and range in composition from hydrogen (atomic number Z = 1) to a predicted atomic number of Z = 114. This composition also includes small percentages of electrons, positrons, and possibly antiprotons. Cosmic-ray astronomy provides information on the origin of the elements (nucleosynthetic processes) and the physics of particles at ultrahigh-energy levels. Such information addresses astrophysical questions concerning the nature of stellar explosions and the effects of cosmic rays on star formation and galactic structure and stability.

Optical Astronomy (Hubble Space Telescope)

Astronomical work in a number of areas has greatly benefited from large, high-resolution optical systems that have operated or are now operating outside Earth's atmosphere. Some of these interesting areas include investigation of the interstellar medium, detailed study of quasars and black holes, observation of binary X-ray sources and accretion disks, extragalactic astronomy, and observational cosmology. The Hubble Space Telescope (HST) constitutes the very heart of NASA's space-borne ultraviolet/optical astronomy program at the beginning of the 21st century. Launched in 1990 and repaired and refurbished in orbit by space shuttle crews, the HST's continued ability to cover a wide range of wavelengths, to provide fine angular resolution, and to detect faint sources makes it one of the most powerful and important astronomical instruments ever built.

Ultraviolet Astronomy

Another interesting area of astrophysics involves the extreme ultraviolet (EUV) region of the electromagnetic spectrum. The interstellar medium is highly absorbent at EUV wavelengths (100 to 1,000 angstroms [A°]). EUV data gathered from space-based instruments, such as those placed on NASA's Extreme Ultraviolet Explorer (EUVE), are being used to confirm and refine contemporary theories of the late stages of stellar evolution, to analyze the effects of EUV radiation on the interstellar medium, and to map the distribution of matter in our solar neighborhood.

Infrared Astronomy

Infrared (IR) astronomy involves studies of the electromagnetic (EM) spectrum from 1 to 100 micrometers wavelength, while radio astronomy involves wavelengths greater than 100 micrometers. (A micrometer is one millionth [10^{-6}] of a meter.) Infrared radiation is emitted by all classes of "cool" objects (stars, planets, ionized gas and dust regions, and galaxies) and the cosmic background radiation. Most emissions from objects with temperatures ranging from about 3 to

2,000 kelvins are in the infrared region of the spectrum. In order of decreasing wavelength, the sources of infrared and microwave (radio) radiation are:

(1) Galactic synchrotron radiation (The radiation emitted by charged particles spiraling around magnetic lines of force in an applied magnetic fieled is called ***synchrotron radiation***)

(2) Galactic thermal bremsstrahlung radiation in regions of ionized hydrogen.

(3) The cosmic background radiation.

(4) 15-kelvins cool galactic dust and 100-kelvins stellar- heated galactic dust.

(5) Infrared galaxies and primeval galaxies.

(6) 300-kelvins interplanetary dust.

(7) 3,000- kelvins starlight.

Advanced space-based IR observatories, such as NASA's Spitzer Space Telescope, are revolutionizing how astrophysicists and astronomers perceive the universe. The Spitzer Space Telescope is the fourth and final element in NASA's family of orbiting great observatories, which includes the Hubble Space Telescope, the Compton Gamma Ray Observatory, and the Chandra X-Ray Observatory. These great space-based observatories give astrophysicists and astronomers an orbiting "toolbox" that provides multi-wavelength studies of the universe.

The Appropriate Radiation Laws

All macroscopic objects; fires, ice cubes, people, stars emit radiation at all times, regardless of their size, shape, or chemical composition. They radiate mainly because the microscopic charged particles they are made up of are in constantly varying random motion, and whenever charges change their state of motion, electromagnetic radiation is emitted. The temperature of an object is a direct measure of the amount of microscopic motion within it. The hotter the object that is, the higher its temperature, the faster its constituent particles move and the more energy they radiate.

THE BLACKBODY SPECTRUM

Intensity is a term often used to specify the amount or strength of radiation at any point in space. Like frequency and wavelength, intensity is a basic property of radiation. No natural object emits all its radiation at just one frequency. Instead, the energy is generally spread out over a range of frequencies. By studying the way in which the intensity of this radiation is distributed across the electromagnetic spectrum, we can learn much about the object's properties.

Figure 1 illustrates schematically the distribution of radiation emitted by any object. The curve peaks at a single, well-defined frequency and falls off to lesser values above and below that frequency. Note that the curve is not shaped like a symmetrical bell that declines evenly on either side of the peak. The intensity falls off more slowly from the peak to lower frequencies than it does from the peak to high frequencies. This overall shape is characteristic of the radiation emitted by any object, regardless of its size, shape, composition, or temperature.

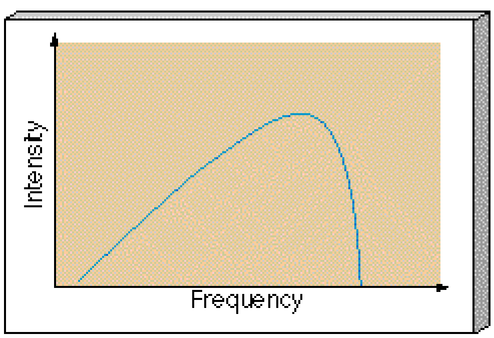


Figure 1: *The blackbody, or Planck, curve represents the distribution of the intensity of radiation emitted by any heated object.*

The curve drawn in Figure 1 is the radiation-distribution curve for a mathematical idealization known as a blackbody, an object that absorbs all radiation falling on it. In a steady state, a blackbody must re emit the same amount of energy as it absorbs. The blackbody curve shown in the Figure 1 describes the distribution of that re emitted radiation. (The curve is also known as the Planck curve, after Max Planck, whose mathematical analysis of such thermal emission in 1900 played a key role in modern physics.) No real object absorbs and radiates as a perfect blackbody. However, in many cases, the blackbody curve is a very good approximation to reality, and the properties of blackbodies provide important insights into the behavior of real objects.

THE RADIATION LAWS

The blackbody curve shifts toward higher frequencies (shorter wavelengths) and greater intensities as an object's temperature increases. Even so, the shape of the curve remains the same. This shifting of radiation's peak frequency with temperature is familiar to us all: very hot glowing objects, such as toaster filaments or stars, emit visible light. Cooler objects, such as warm rocks or household radiators, produce invisible radiation warm to the touch but not glowing hot to the eye. These latter objects emit most of their radiation in the lower-frequency infrared part of the electromagnetic spectrum.

Imagine a piece of metal placed in a hot furnace. At first, the metal becomes warm, although its visual appearance doesn't change. As it heats up, it begins to glow dull red, then orange, brilliant yellow, and finally white. How do we explain this? As illustrated in Figure 2, when the metal is at room temperature (300 K), it emits only invisible infrared radiation. As the metal becomes hotter, the peak of its blackbody curve shifts toward higher frequencies. At 1000 K, for instance, most of the emitted radiation is still infrared, but now there is also a small amount of visible (dull red) radiation being emitted (note in Figure 2 that the high-frequency portion of the 1000 K curve just overlaps the visible region of the graph).

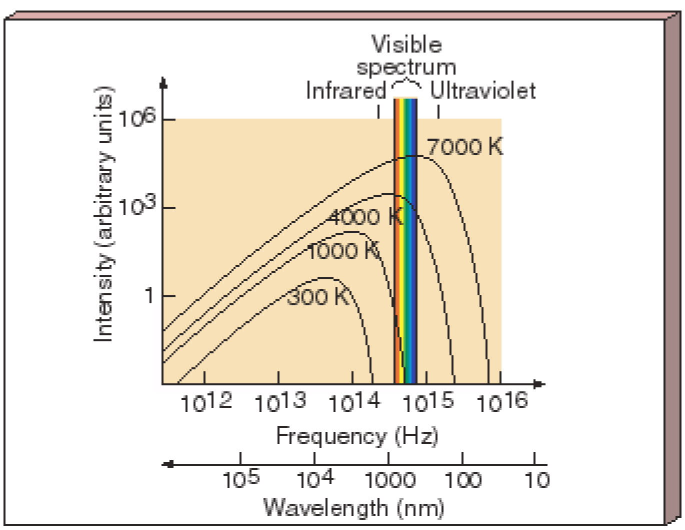


Figure 2 *As an object is heated the radiation it emits peaks at higher and higher frequencies. Shown here are curves corresponding to temperatures of 300 K (room temperature), 1000 K (beginning to glow deep red), 4000 K (red hot), and 7000 K (white hot).*

Blackbody Radiation

A blackbody is a theoretical idea that closely approximates many real objects in thermodynamic equilibrium. (**We say that an object is in thermodynamic equilibrium with its surroundings when energy is freely interchanged and a steady state is reached in which there is no net energy flow. That is, energy flows in and out at the same rate**.) A blackbody is an object that absorbs all of the radiation that strikes it.

A blackbody can also emit radiation. In fact, if a blackbody is to maintain a constant temperature, it must radiate energy at the same rate that it absorbs energy. If it radiates less energy than it absorbs, it will heat up. If it radiates more energy than it absorbs, then it will cool. However, this does not mean that the spectrum of emitted radiation must match the spectrum of absorbed radiation. Only the total energies must balance. The spectrum of emitted radiation is determined by the temperature of the blackbody.

Wien's displacement law

Wien's displacement law, named after Wilhelm Wien was derived in the year 1893 which states that black body radiation has different peaks of temperature at wavelengths that are inversely proportional to temperatures.

The relationship between the wavelength at which the peak occurs, λ_{max}, and temperature, T, is very simple. It is given by:

$$\lambda_{max} T = 2.90 \times 10^{-1} \text{ cm K} = 2.90 \times 10^{6} \text{ nm K} \quad \ldots\ldots 1$$

In this law, we must use temperature on an absolute (Kelvin) scale. (The temperature on the Kelvin scale is the temperature on the Celsius scale plus 273.1.)

Example 1 Using Wien's displacement law (a) Find the temperature of an object whose blackbody spectrum peaks in the middle of the visible part of the spectrum, λ = 550 nm. (b) The Earth has an average temperature of about 300 K. At what wavelength does the Earth's blackbody spectrum peak?

SOLUTION

(a) Given the wavelength, we solve equation (1) for the temperature:

$$T = \frac{2.9 \times 10^{6} \text{nm K}}{550 \text{ nm}} = 5270 \text{ K}$$

This is close to the temperature of the Sun.

(b) Given the temperature, we solve equation (1) for the wavelength:

$$\lambda = \frac{2.9 \times 10^6 \text{ nm K}}{300 \text{ K}}$$
$$= 1 \times 10^4 \text{ nm}$$
$$= 10 \times 10^{-6} \text{ m}$$
$$= 10 \ \mu\text{m}$$

This is in the infrared part of the spectrum. Even though the Earth is giving off radiation, we don't see it glowing in the visible part of the spectrum. Similarly, objects around us that are at essentially the same temperature as the Earth give off most of their radiation in the infrared part of the spectrum, with very little visible light. The visible light that we see from surrounding objects is partially reflected sunlight or artificial light.

We could have solved (b) by scaling a known result, such as the answer in (a):

$$\frac{\lambda_{\text{Earth}}}{\lambda_A} = \frac{T_A}{T_{\text{Earth}}}$$
$$\lambda_{\text{Earth}} = \left(\frac{T_A}{T_{\text{Earth}}}\right)\lambda_A$$
$$= \left(\frac{5270 \text{ K}}{300 \text{ K}}\right)(550 \text{ nm})$$
$$= 10 \times 10^3 \text{ nm}$$
$$= 10 \ \mu\text{m}$$

Scaling results can be useful because they show how different physical parameters are related to each other. It also provides us with a way of using an equation even if we don't remember the constants.

The Stefan–Boltzmann law

Suppose we are interested in the total energy given off by a blackbody (per unit time per unit surface area) over the whole electromagnetic spectrum. We must add the contributions at all wavelengths. This amounts to taking an integral over blackbody curves. Since a hotter blackbody gives off more energy at all wavelengths than a cooler one, and is particularly dominant at shorter wavelengths, we would expect a hotter blackbody to give off much more energy than a cooler one. Indeed, this is the case. ***The total energy per unit time, per unit surface area, E, given off by a blackbody is proportional to the fourth power of the temperature. That is***

$$E = \sigma T^4 \qquad \text{……………………………..2}$$

This relationship is called **the Stefan–Boltzmann law**. The constant of proportionality, **σ**, is called the Stefan–Boltzmann constant. It has a value of 5.7×10^{-5} erg / (cm^2 K^4 s). This law was first determined experimentally, but it can also be derived theoretically. The $\mathbf{T^4}$ dependence means that ***E*** depends strongly on **T**. If we double the temperature of an object, the rate at which it gives off energy goes up by a factor of 16. If we change the temperature by a factor of ten (say from 300 K to 3000 K), the energy radiated goes up by a factor of 10^4. For a star, we are interested in the total luminosity. **The luminosity** is the total energy per second (i.e. the power) given off by the star. The quantity $\boldsymbol{\sigma}$ $\mathbf{T^4}$ is only the energy per second per unit surface area. Therefore, to obtain the luminosity, we must multiply it by the surface area. If the star is a sphere with radius ***R***, the surface area is ($4\pi R^2$), so the luminosity is

$$L = (4\pi R^2)(\sigma T^4) \qquad \text{………………………..3}$$

Example 2 The surface temperature of the Sun is about 5800 K and its radius is 7×10^5 km (7×10^{10} cm). What is the luminosity of the Sun?

SOLUTION

We use equation (3) to find the luminosity:

$$L = 4\pi(7 \times 10^{10}\ \text{cm})^2\ [5.7 \times 10^{-5}\ \text{erg/(cm}^2\ \text{K}^4\ \text{s)}] \times (5.8 \times 10^3\ \text{K})^4$$
$$= 4 \times 10^{33}\ \text{erg/s}.$$

This quantity is called the solar luminosity, and serves as a convenient unit for expressing the luminosities of other stars.

The Planck's law

The primary law governing blackbody radiation is the Planck Radiation Law. This law give us the relation between the intensity of radiation emitted by unit surface area from the blackbody into a fixed direction as a function of wavelength for a fixed temperature. The Planck's law can be expressed through the following equation.

$$E(\lambda, T) = \frac{2hc^2}{\lambda^5}[\exp(\frac{hc}{\lambda kT}) - 1]^{-1} \qquad \text{…………………….4}$$

h = the Planck constant = 6.63×10^{-34} J s ,
c = the speed of light ≈ 3×10^8 m s^{-1} ,
k = the Boltzmann constant = 1.38×10^{-23} J K^{-1}

The behavior of equation (4) is illustrated in the Figure 2. The Planck's law gives us a distribution of intensity of radiation that peaks (max. point) at a certain wavelength. From the Figure 2 we notice that : the peak shifts to shorter wavelengths for higher temperatures, and the area under the curve grows rapidly with increasing temperature.

Example:

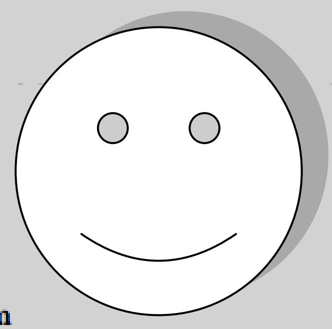
How many visible photon λ =5000 A° are emitted each second by a 100 watt light blue?

Answer:

A photon is characterized by either a wavelength, denoted by λ or equivalently an energy, denoted by E. There is an inverse relationship between the energy of a photon (E) and the wavelength of the light (λ) given by the equation:

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

where h is Planck's constant and c is the speed of light.



h = 6.626 × 10 $^{-34}$ joule·s

c = 2.998 × 10^{8} m/s

By multiplying to get a single expression, hc = 1.99 × 10^{-25} joules-m

The above inverse relationship means that light consisting of high energy photons (such as "blue" light) has a short wavelength. Light consisting of low energy photons (such as "red" light) has a long wavelength.

When dealing with "particles" such as photons or electrons, a commonly used unit of energy is the electron-volt (eV) rather than the joule (J). An electron volt is the energy required to raise an electron through 1 volt, thus a photon with an energy of 1 eV = 1.602 × 10^{-19} J.

Therefore, we can rewrite the above constant for hc in terms of eV:

hc = (1.99 × 10^{-25} joules-m) × (1ev/1.602 × 10^{-19} joules) = 1.24 × 10^{-6} eV-m

1.24 × 10^{-6} eV.m × 10^{10} = 12400 eV . A°

For Blue Light

The energy of each photon is E_{photon} = hc / λ =12400 eV.A° / 4000 A° = 2.48 eV = 3.97x10^{-12} erg

1 watt = 10^{7} erg. s^{-1}

Total energy = 100 watt = 10^9 erg. s^{-1}

∴ The number of the visible photons per second = total energy / E_{photon}

$=10^9 / 3.97 \times 10^{-12} = 2.52 \times 10^{20}$.

Exercise: Calculate the energy of photons, in eV and in watts, for wavelengths of 100, 1000, 10 000, and 100 000 angstroms. How do these energies compare with each other? How do they compare to a 100-watt light bulb? How many 5000! photons would you need to make the equivalent of a 100-watt light bulb?

Stellar magnitude

Greek astronomers were thought to be the first to classify stars by their brightness. Around 129 B. C. Hipparchus ranked the 1,080 stars in his catalogue in a simple way; the brightest were termed “first magnitude,” the next brightest were “second magnitude” through to the “sixth magnitude.” It was a rough estimation of star brightness. Our knowledge of Hipparchus is via Ptolemy, who, in the 2^{nd} century A. D., utilized a similar stellar brightness scheme in his Almagest. Of the 1, 028 stars in his catalogue, 156 were given the descriptions of a little more or a little less than the integer values, but his precision was off by ± 0.6 magnitudes. The Persian astronomer Abu’l-Husayn al-Sufi re-estimated the Almagest’s magnitude values in the tenth century. Tycho Brahe (1546-1601) added to the list of magnitudes, but with no precision improvement.

The introduction of the telescope allowed astronomers to see fainter stars. Galileo (1564-1642) estimated he could see magnitude +12 stars. However, it wasn’t until William Herschel (1738-1822) that magnitude accuracy reached ± 0.2 magnitudes. He felt that brightness was solely based on stellar distance, and (unsuccessfully) attempted to delineate the Milky Way based on the distribution of stars of different magnitudes. To standardize magnitudes, Norman Pogson (1829-1891) proposed that a brightness ratio of 100:1 be equivalent to five magnitudes. In other words, a difference of one magnitude implied a brightness difference of **2.512**. This Pogson ratio is now the accepted magnitude standard. Today, the eyeball has been supplanted by CCD photometry for very accurate magnitude estimation.

Old astronomers used magnitude scale as follow:

Star magnitude	1	2	3	4	5	6
Brightness Ratio	100					1.00

* The brightest stars has magnitude = 1
* The faintest stars that could be observed by the naked eye = 6.
* stars of magnitudes larger than 6 needs telescopes
* the larger the number the fainter the star. Therefore we can say that the magnitude scale is a measure of how the star is faint.

* Arabs used this magnitude scale to measure and test the eye efficiency. Who can not observe star of magnitude 6 his eye measure is 5/6 or lower depending on the stellar magnitudes that he can see. Who can see the stars with m = 6 is the best eye visibility.

* We can reorganize the magnitude scale as follows:

Object	Magnitude
Sun	-26.5
Full Moon	-12.5
Venus (at brightest)	-4
Mars, Jupiter (at brightest)	-2
Sirius	-1.5
Aldebran, Altair	1.0
Naked eye limit	6.5
5-m optical telescope	20
Hubble Space telescope	30

Brightness b: is the flux of radiation received (arrived at Earth) from the star.

The relation between the magnitudes and the brightness are established In the late 18th and 19th centuries, several astronomers performed experiments to see how the magnitude scale was related to the amount of energy received. It appeared that a given difference in magnitude, at any point in the magnitude scale, corresponded to a ratio of the brightnesses. In 1856 Pogson proposed that the value of the ratio, corresponding to a magnitude difference of **five**, should be **100**. Thus, the ratio of two stellar brightnesses, B1 and B2, can be related to their magnitudes, m1 and m2, by the equation

$$\frac{B_1}{B_2} = 2{\cdot}512^{-(m_1-m_2)} \qquad \ldots\ldots\ldots\ldots\ldots\ldots\ldots 1$$

Since $(2{\cdot}512)^5$ equals 100. This is known as **Pogson's equation**. The negative sign before the bracketed exponent reflects the fact that magnitude values increase as the brightness falls. By taking logarithms of equation above, we obtain

$$\log_{10}\left(\frac{B_1}{B_2}\right) = -(m_1 - m_2)\log_{10}(2{\cdot}512) = -0{\cdot}4(m_1 - m_2)$$

$$m_1 - m_2 = -2{\cdot}5\log_{10}\left(\frac{B_1}{B_2}\right) \qquad \ldots\ldots\ldots\ldots\ldots\ldots\ldots 2$$

More generally, Pogson's equation in the style of equation (1) can be presented in a simplified form as

$$m = k - 2.5 \log_{10} B \quad \ldots\ldots\ldots\ldots\ldots\ldots\ldots\ldots 3$$

Where *m* is the magnitude of the star, *B* its apparent brightness and *k* some constant. The value of *k* is chosen conveniently by assigning a magnitude to one particular star such as α Lyr, or set of stars, thus fixing the zero point to that magnitude scale. It should also be noted that the numerical coefficient of 2·5 in equation (2) is exact and is not a rounded value of 2·512 from equation (1).

Luminosity and Absolute magnitude

Luninosity: luminosity is the total amount of energy emitted by a star, galaxy, or other astronomical object per unit time.

We define apparent magnitude as a logarithmic luminosity ratio of a body to some standard:

$$m = -2.5\log(L/L_0) \quad \ldots\ldots\ldots\ldots\ldots 4$$

Where *m* is the apparent magnitude, *L* is the luminosity of a body as determined through a V filter, and L_0 is the standard's luminosity, the luminosity of Vega. Vega has been assigned an $m = 0$ although its actual magnitude is 0.03. It follows that

$$m_1 - m_2 = -2.5\log(L_1/L_2) \quad \ldots\ldots\ldots\ldots\ldots\ldots\ldots 5$$

Equation 1 can be rewritten as:

$$L = L_0 x 10^{-m/2.5} \quad \ldots\ldots\ldots\ldots\ldots\ldots\ldots\ldots\ldots\ldots\ldots\ldots\ldots\ldots 6$$

EXAMPLE 1. The Sun is about 480,000 times more luminous than the full Moon. What is the difference in their apparent magnitude?

$L_{moon} = 1$, $L_{Sun} = 480{,}000$ $\quad \Delta m = -2.5\log(\frac{L_{Sun}}{L_{moon}}) = -2.5\log(\frac{480,000}{1}) = -14.2$ magnitude.

The Sun's apparent magnitude is -26.8, while the full moon's is -12.6.

EXAMPLE 2. The individual apparent magnitudes of two binary stars are +2 and +4. What is the combined apparent magnitude of the binary system?

It's obvious we cannot simply add or subtract the two magnitudes. We must revert to their original luminosities using equation 3, which are additive, then convert the sum to magnitude using Equation 1.

$$L_{binary}/L_0 = L_{+2}/L_0 + L_{+4}/L_0 = 10^{-4/2.5} + 10^{-2/2.5} = 0.1836 \text{ from Equation}$$

$$m_{binary} = -2.5\log(L_{binary}/L_0) = -2.5\log(0.1836) = 1.840 \text{ magnitude}$$

EXAMPLE 3. Sirius is 8.6 light years distant, with an apparent magnitude of -1.5. What would the apparent magnitude of the Sun be at Sirius' distance?

There are 63,241 AU in a ly so Sirius is 543,873 AU distant. The Sun would be $1/(543{,}873)^2$ fainter at Sirius' distance. At 8.6 ly, the Sun's magnitude would change by

$$\Delta m_{Sun} = -2.5\log(1/(543{,}873)^2) = 28.7 \text{ magnitudes fainter}$$

So the Sun's apparent magnitude would be -26.8 + 28.7 = +1.9 at Sirius' distance, a star slightly brighter than Polaris.

The inverse square law for light:

The brightness (***B***) of an object decays with distance (***d***) as follows:

$$B \propto \frac{1}{d^2} \quad \ldots\ldots\ldots\ldots\ldots\ldots\ldots 7$$

As the star distance increases its brightness decreases as the square of the star distance from us. This is a general physical law.

Absolute magnitude

Since stellar brightness depends on distance, astronomers utilize magnitudes at a standard distance, 10 parsecs (32.62 ly), to compare stellar intrinsic luminosity. We already determined in Example 3 that if the Sun was at Sirius's distance, it would appear 28.7 magnitudes fainter.

We'll relate three quantities: apparent magnitude, m; absolute magnitude, M; and stellar distance in parsecs, d. (1 parsec = 3.26 light years.)

$$m - M = 5 \log (d) - 5 \quad \text{.....................7}$$

Equation 4 can be rewritten as follows:

$$d = 10^{\frac{m-M+5}{5}} \quad \text{........................8}$$

The expression (m – M) is called the distance modulus. The distance modulus is negative if a star is closer than 10 parsecs.

EXAMPLE 4. Find the absolute magnitude of the Sun with m = -26.8.

We need the Sun's distance to be in parsecs. There are 206,265 AU in a parsec, so d = 1/206,265. Using Equation 4,

$$M = m - 5 \log (d) + 5 = -26.8 - 5 \log (1/206{,}265) + 5 = 4.77$$

EXAMPLE 5. A star's m = 8 and M = -2. Find the star's distance.

Notice that the distance modulus is 8 – (-2) = 10. This indicates the star is more distant than 10 parsecs. Using Equation 5,

$$d = 10^{\frac{m-M+5}{5}} = 10^{\frac{8-(-2)+5}{5}} = 10^{3} = 1{,}000 \text{ parsecs.}$$

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Luminosity, Radius, and Temperature

If a star is considered simply as a spherical source radiating as a black body, according to its surface temperature and its surface area. This total output is referred to as the stellar luminosity, *L*, and may be expressed as

$$L = 4\pi R^2 \sigma T^4 \quad \text{W}$$

This seems complicated, but if we express Luminosity, Radius, and Temperature in terms of the Sun, we get a much simpler form:

$$\frac{L_{star}}{L_{sun}} = \left(\frac{R_{star}}{R_{sun}}\right)^2 \left(\frac{T_{star}}{T_{sun}}\right)^4 \qquad \ldots\ldots\ldots\ldots\ldots\ldots\ldots 9$$

Example 6. Suppose we want to find the luminosity of a star 10 times the Sun's radius but only half as hot. How luminous is it?

$$\frac{L_{star}}{L_{sun}} = \left(\frac{10}{1}\right)^2 \left(\frac{1}{2}\right)^4 = 6.25$$ the star has 6.25 times the Sun's luminosity

Bolometric magnitude (M_{blo}) :

If we were able to measure the radiation at all wavelength , we would get the Bolometric magnitude $\boldsymbol{M_{blo}}$ of star. This magnitude is given by

$$M_{blo} = m_v - BC$$

Where $\boldsymbol{m_v}$ is the visual magnitude (the magnitude corresponding to the sensitivity of the eye which occurred at wavelength of 550 nm. BC is the bolometric correction

Example : The distance of a star is 100Pc and its apparent magnitude is 6. What is its absolute magnitude?

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

$6 - M = 5\log(100) - 5$

$M = 1$

*Example :*If absolute magnitude of a star is 1.4 and the bolometric correction is -0.6. What is the bolometric of this star?

Answer: $M_{bol} = M + BC = 1.4 - 0.6 = 0.8$

Bayer designation: is a stellar designation in which a specific star is identified by a Greek letter, followed by the genitive form of its parent constellation's Latin name. The original list of Bayer designations contained 1,564 stars.

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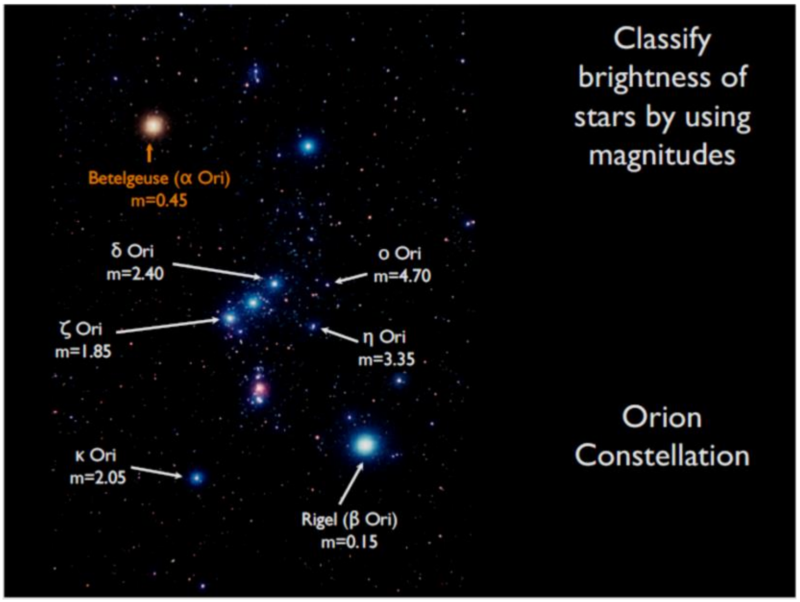

Letter	Name	Letter	Name
α	alpha,	ν	nu,
Β	beta,	ξ	xi,
γ	gamma,	ο	omicron,
δ	delta,	π	pi,
ε	epsilon,	ρ	rho,
ζ	zeta,	σ	sigma,
η	eta,	τ	tau,
θ	theta,	υ	upsilon,
ι	iota,	φ	phi,
κ	kappa,	χ	chi,
λ	lambda,	ψ	psi,
μ	mu,	ω	omega,

Stellar Evolution

Stellar evolution is a description of the way that stars change with time. On human timescales, most stars do not appear to change at all, but if we were to look for billions of years, we would see how stars are born, how they age, and finally how they die.

The primary factor determining how a star evolves is its mass as it reaches the main sequence. The following is a brief outline tracing the evolution of a low-mass and a high-mass star.

The life of a star

Stars are born out of the gravitational collapse of cool, dense molecular clouds.

Molecular Clouds: Dust and gas primarily in the form of hydrogen molecules are the main constituents of the coldest, densest clouds in the interstellar medium. These molecular clouds (the largest of which are known as Giant Molecular Clouds) have typical temperatures of around 10 Kelvin and densities upward of 102 particles/cm3, masses ranging from a few to over a million solar masses and diameters from 20 to 200 parsecs. Star formation takes place exclusively within molecular clouds and observations have shown that they are located primarily in the disk of spiral galaxies and the active regions of irregular galaxies.

As the cloud collapses, it fragments into smaller regions, which themselves contract to form stellar cores.

The formation of stars begins with the collapse and fragmentation of molecular clouds into very dense clumps. These clumps initially contain ~0.01 solar masses of material, but increase in mass as surrounding material is accumulated through accretion. The temperature of the material also increases while the area over which it is spread decreases as gravitational contraction continues, forming a more stellar-like object in the process. During this time, and up until hydrogen burning begins and it joins the main sequence, the object is known as a **protostar**.

These protostars rotate faster and increase in temperature as they condense, and are surrounded by a protoplanetary disk out of which planets may later form.

The central temperature of the contracting protostar increases to the point where nuclear reactions begin. At this point, hydrogen is converted into helium in the core and the star is born onto the main sequence. For about 90% of its life, the star will continue to burn hydrogen into helium and will remain a main sequence star.

Once the hydrogen in the core has all been burned to helium, energy generation stops and the core begins to contract. This raises the internal temperature of the star and ignites a shell of hydrogen burning around the inert core. Meanwhile, the helium core continues to contract and increase in temperature, which leads to an increased energy generation rate in the hydrogen shell. This causes the star to expand enormously and increase in luminosity – the star becomes a red giant.

Eventually, the core reaches temperatures high enough to burn helium into carbon. If the mass of the star is less than about 2.2 solar masses, the entire core ignites suddenly in a helium core flash. If the star is more massive than this, the ignition of the core is more gentle. At the same time, the star continues to burn hydrogen in a shell around the core.

The star burns helium into carbon in its core for a much shorter time than it burned hydrogen. Once the helium has all been converted, the inert carbon core begins to contract and increase in temperature. This ignites a helium burning shell just above the core, which in turn is surrounded by a hydrogen burning shell.

What happens next depends on the mass of the star

Stars less than 8 solar masses

The inert carbon core continues to contract but never reaches temperatures sufficient to initiate carbon burning. However, the existence of two burning shells leads to a thermally unstable situation in which hydrogen and helium burning occur out of phase with each other. This thermal pulsing is characteristic of asymptotic giant branch stars.

The carbon core continues to contract until it is supported by electron degeneracy pressure. No further contraction is possible (the core is now supported by the pressure of electrons, not gas pressure), and the core has formed a white dwarf. Meanwhile, each thermal pulse causes the outer layers of the star to expand, resulting in a period of mass loss. Eventually, the outer layers of the star are ejected completely and ionised by the white dwarf to form a planetary nebula.

Stars greater than 8 solar masses

The contracting core will reach the temperature for carbon ignition, and begin to burn to neon. This process of core burning followed by core contraction and shell burning, is repeated in a series of nuclear reactions producing successively heavier elements until iron is formed in the core.

Iron cannot be burned to heavier elements as this reaction does not generate energy – it requires an input of energy to proceed. The star has therefore finally run out of fuel and collapses under its own gravity.

The mass of the core of the star dictates what happens next. If the core has a mass less than about 3 times that of our Sun, the collapse of the core may be halted by the pressure of neutrons (this is an even more extreme state than the electron pressure that supports white dwarfs!). In this case, the core becomes a neutron star. The sudden halt in the contraction of the core produces a shock wave which propagates back out through the outer layers of the star, blowing it apart in a core-collapse supernova explosion. If the core has a mass greater than about 3 solar masses, even neutron pressure is not sufficient to withstand gravity, and it will collapse further into a stellar black hole.

The ejected gas expands into the interstellar medium, enriching it with all the elements synthesised during the star’s lifetime and in the explosion itself. These supernova remnants are the chemical distribution centres of the Universe.

An important tool in the study of stellar evolution is the Hertzsprung-Russell diagram (HR diagram), which plots the absolute magnitudes of stars against their spectral type (or alternatively, stellar luminosity versus effective temperature). As a star evolves, it moves to specific regions in the HR diagram, following a characteristic path that depends on the star’s mass and chemical composition.

Hertzsprung-Russell Diagram

The Hertzsprung-Russell diagram (HR diagram) is one of the most important tools in the study of stellar evolution. Developed independently in the early 1900s by Ejnar Hertzsprung and Henry Norris Russell, it plots the temperature of stars against their luminosity (the theoretical HR diagram), or the colour of stars (or spectral type) against their absolute magnitude (the observational HR diagram, also known as a colour-magnitude diagram).

Depending on its initial mass, every star goes through specific evolutionary stages dictated by its internal structure and how it produces energy. Each of these stages corresponds to a change in the temperature and luminosity of the star, which can be seen to move to different regions on the HR diagram as it evolves. This reveals the true power of the HR diagram – astronomers can know a star's internal structure and evolutionary stage simply by determining its position in the diagram.

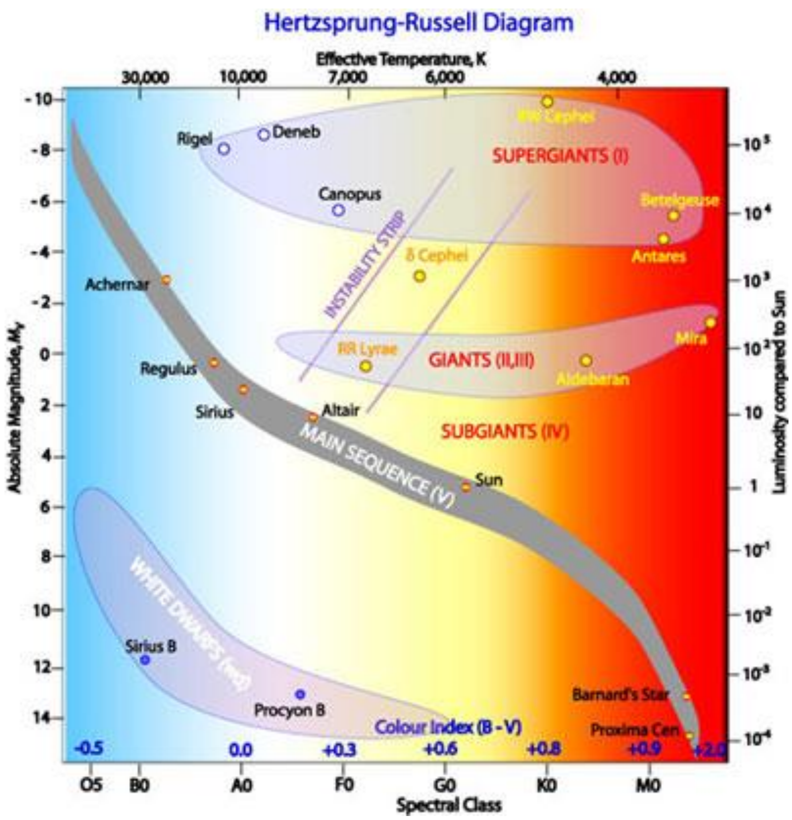

Figure: The Hertzsprung-Russell diagram the various stages of stellar evolution. By far the most prominent feature is the main sequence (grey), which runs from the upper left (hot, luminous stars) to the bottom right (cool, faint stars) of the diagram. The giant branch and supergiant stars lie above the main sequence, and white dwarfs are found below it.

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There are 3 main regions (or evolutionary stages) of the HR diagram:

1. The **main sequence** stretching from the upper left (hot, luminous stars) to the bottom right (cool, faint stars) dominates the HR diagram. It is here that stars spend about 90% of their lives burning **hydrogen** into **helium** in their cores. Main sequence stars have a **Morgan-Keenan luminosity class** labelled **V**.
2. **red giant** and **supergiant** stars (luminosity classes **I** through **III**) occupy the region above the main sequence. They have low surface temperatures and high **luminosities** which, according to the Stefan-Boltzmann law, means they also have large radii. Stars enter this evolutionary stage once they have exhausted the hydrogen fuel in their cores and have started to burn helium and other heavier elements.
3. **white dwarf** stars (luminosity class **D**) are the final evolutionary stage of low to intermediate mass stars, and are found in the bottom left of the HR diagram. These stars are very hot but have low luminosities due to their small size.

The **Sun** is found on the main sequence with a luminosity of 1 and a temperature of around 5,400 **Kelvin**.
Astronomers generally use the HR diagram to either summarise the evolution of stars, or to investigate the properties of a collection of stars. In particular, by plotting a HR diagram for either a globular or open cluster of stars, astronomers can estimate the age of the cluster from where stars appear to turn off the main sequence (see the entry on main sequence for how this works).

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Originally, stars were assigned a type A to Q based on the strength of the hydrogen lines present in their spectra. However, it was later realised that there was significant overlap between the types, and some of the letters were dropped. Continuity of other spectral features was also improved if B came before A and O came before B, with the end result, the spectral sequence: OBAFGKM. This sequence is ordered from the hottest to the coolest stars, and is often remembered by the mnemonic 'Oh Be A Fine Girl/Guy, Kiss Me'.

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few visible absorption lines, weak **Balmer lines**, ionised **helium** lines	**neutral hydrogen** lines, more prominent Balmer lines	strongest Balmer lines, other strong lines	weaker Balmer lines, many lines including neutral **metals**	Balmer lines weaker still, dominant ionised calcium lines	neutral metal lines most prominent	strong neutral metal lines and molecular bands

Unfortunately, proper classification of a stellar spectrum is not quite this simple. Within each spectral type there are significant variations in the strengths of the absorption lines, and each type has been subdivided into 10 sub-classes numbered 0 to 9. In addition, stars of a particular spectral type can differ widely in luminosity and must also be assigned a luminosity class. This distinguishes main sequence stars (dwarf stars) from giant and supergiant stars.

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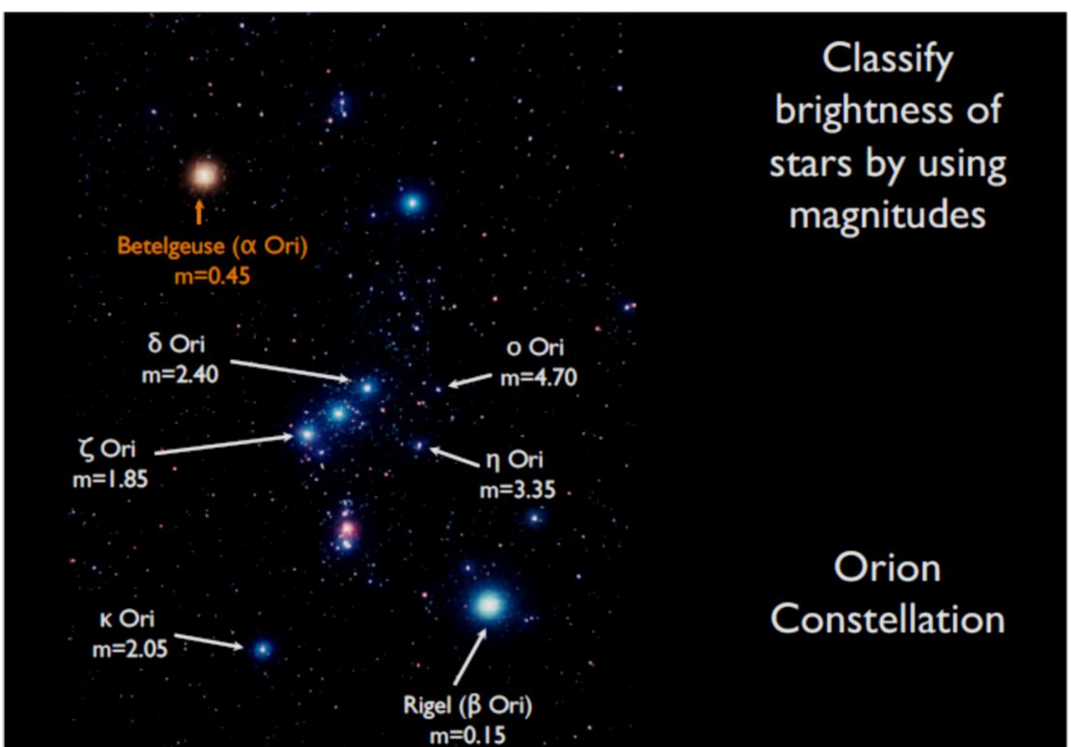

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ζ	zeta,

Letter	Name
ν	nu,
ξ	xi,
ο	omicron,
π	pi,
ρ	rho,
σ	sigma,

η	eta,
θ	theta,
ι	iota,
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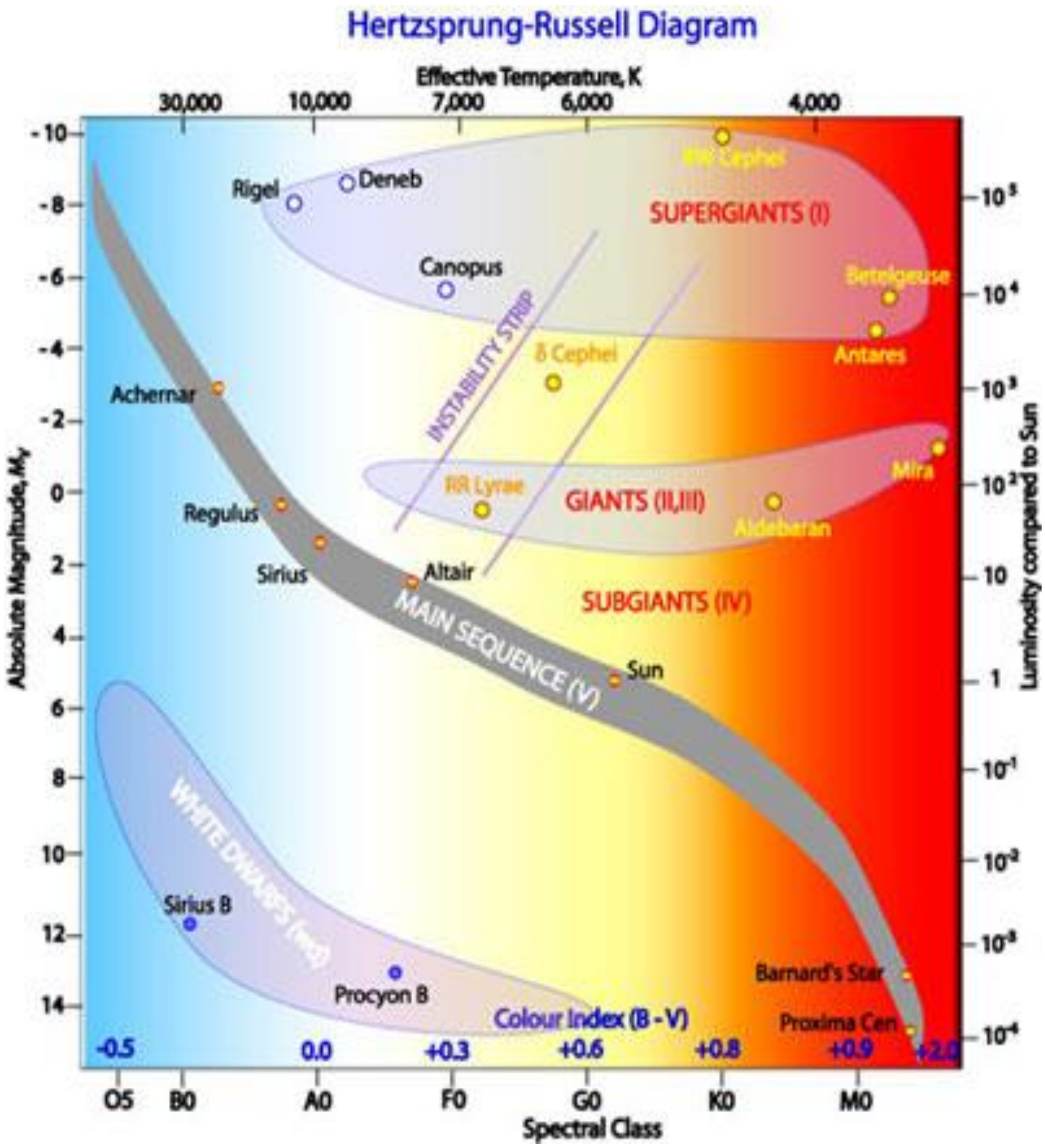

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3.8 STELLAR PARALLAX (TRIGONOMETRIC) AND THE UNITS OF STELLAR DISTANCES

Stellar parallax is a measure of the star's distance. The diameter of the Earth's orbit around the Sun is used as a basic line of such measurements and the corresponding parallax is called the trigonometric parallax. Thus, the trigonometric parallax of a star may be defined as the angle p (in seconds of arc) subtended at the star σ by the mean radius 'a' of the Earth's orbit round the Sun. This is also called heliocentric or annual parallax. A nearby star σ is photographed from the position E of the Earth against the background of the far off stars. Similar observations are also made after six months, from the position E' of the Earth on the other end of the base line. From these observations, the angle subtended by the diameter EE' (or base line) at the star σ is measured, comparing the change in the position of σ during the six months with respect to the fainter far off stars whose shift in position due to the motion of the Earth is negligible. One-half of this angle gives the measure of annual parallax p of the star. The sets of observations at E and E' are repeated when the Earth comes back to this position after a year and the average of a number of measurements is taken as the parallax of the star. As the stars are too far away, the parallax of a star is a very small angle and its value has been found to be always less than 1″ of arc even for the nearest stars.

From Fig. 3.1, we have

$$p_{\text{rad}} = \frac{a}{d}, \text{ where } d = \text{distance of the star.}$$

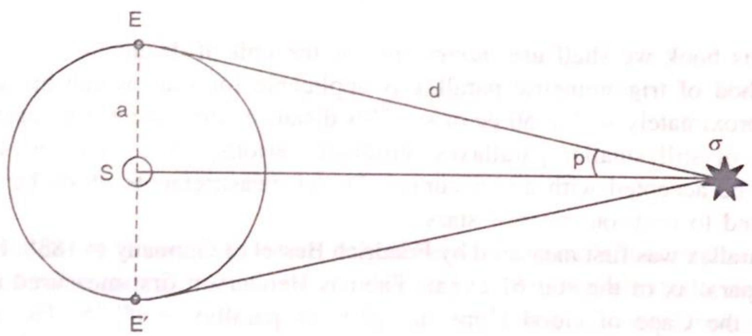


FIGURE 3.1 The annual parallax of a star.

But

$$1_{\text{rad}} = 206{,}265'' \text{ of arc.}$$

Therefore,

$$\frac{p''}{206{,}265} = \frac{a}{d},$$

since the parallax of a star is expressed in seconds of arc, or

$$d = \frac{206{,}265\, a}{p''} \tag{3.12}$$

Planetary nebulae

We have already said that the outer layers of a red giant are held together very weakly. Remember, the gravitational force on a mass m in the outer layer is GmM/R^2, where M is the mass of the star and R is its radius. As the star expands, M stays constant, so the pull on the outer layer falls off as $1/R^2$. Since the outer layer is weakly held, it is subject to being driven away. The actual mechanism for driving material away is still not fully understood. It may involve pressure waves moving radially outward. It may also involve radiation pressure. Photons carry energy and momentum. (Remember, the momentum of a photon of energy E is E/c.) When photons from inside the star strike the gas in the outer layers, and are absorbed, their momentum is also absorbed. By conservation of momentum, the shell must move slightly outward. We do observe shells that are ejected. They are fuzzy in appearance in small telescopes, just like planets; when originally observed, they were called planetary nebulae (Fig. 10.8). (Their name has nothing to do with their properties, but with their appearance as viewed with small telescopes.) From the photograph in Figure below we see that some planetary

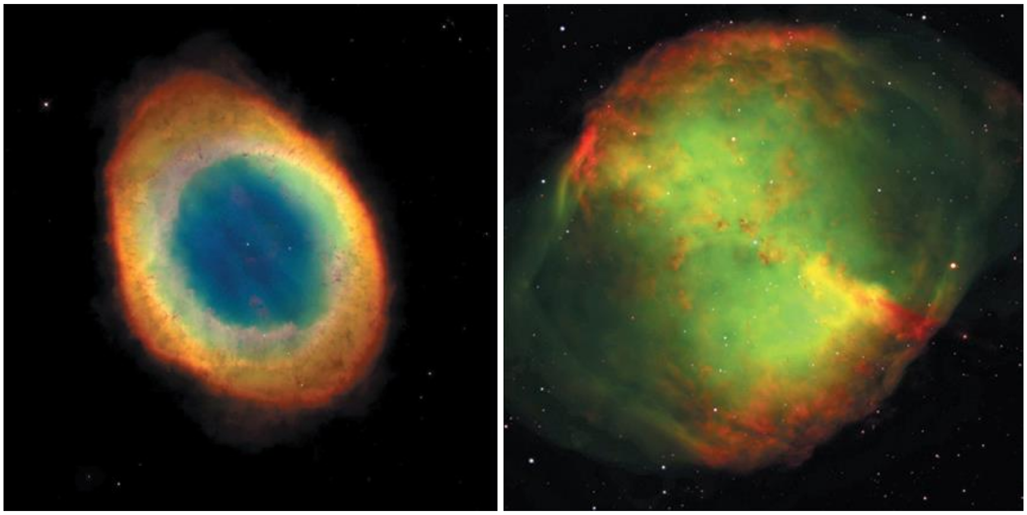


Figure: *Images of planetary nebulae. (a) HST image of the Ring Nebula (M57), in the constellation Lyra. It is at a distance of 1 kpc, and is about 0.3 pc across.This image reveals elongated dark clumps of material at the edge of the nebula. (b) The Dumbbell Nebula (M27), in a ground-based image.This is 300 pc away and is 0.5 pc across, in the constellation of Vulpecula.*

nebulae have a ringlike appearance. However, they are spherical shells. We see them as rings because our line of sight through the edge of the shell passes through more material than the line of sight through the center (Figure below).

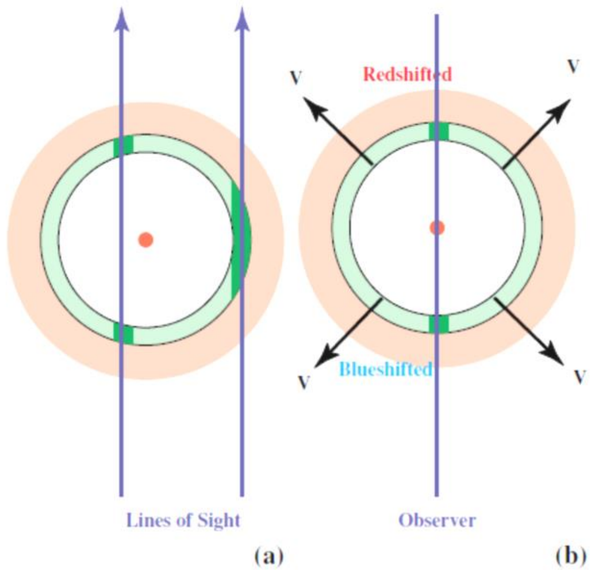


Figure: *Lines of sight through a planetary nebula. (a) Appearance.The shaded regions represent the places where the lines of sight pass though the shell.The line of sight near the edge passes through more material than that through the center.This is responsible for the ringlike appearance. (b) Doppler shifts. Material on the near side is moving toward the observer, producing a blueshift, and material on the far side is moving away, producing a redshift.*

Thus, the center appears to be quite faint. When we look at spectral lines in planetary nebulae, we see two Doppler shifts. One line is redshifted and the other is blueshifted. The blueshifted one comes from the part of the shell that is moving towards us, and the redshifted line comes from the part of the shell that is moving away from us. From the Doppler shifts we find that the shells are expanding at velocities of a few tens of kilometers per second. The physical conditions in planetary nebulae are determined from observations of various spectral lines. Different lines are sensitive to different temperature and density ranges. Information is also obtained from studies of radio waves emitted by the nebulae.

From these studies, we find that masses of planetary nebulae are of the order of 0.1 $M_{\odot}$. The temperatures are about 10^4 K. The mass tells us that up to 10% of the stellar mass is returned to the interstellar medium in the ejection of the nebula. This material will be included in the next generation of stars to form out of the interstellar medium. (This is in addition to mass lost through stellar winds).

………………………………………………………………………………………….

White dwarfs

In ordinary stars the pressure of the gas obeys the equation of state of an ideal gas. In stellar interiors the gas is fully ionized, i. e. it is plasma consisting of ions and free electrons. The partial pressures of the ions and electrons together with the radiation pressure important in hot stars comprise the total pressure balancing gravitation. When the star runs out of its nuclear fuel, the density in the interior increases, but the temperature does not change much. The electrons become degenerate, and the pressure ismainly due to the pressure of the degenerate electron gas, the pressure due to the ions and radiation being negligible. The star becomes a white dwarf.

The radius of a degenerate star is inversely proportional to the cubic root of the mass. Unlike in a normal star the radius decreases as the mass increases. The first white dwarf to be discovered was Sirius B, the companion of Sirius (Fig. below). Its exceptional nature was realized in 1915, when it was discovered that its effective temperature was very high. Since it is faint, this meant that its radius had to be very small, slightly smaller than that of the Earth. The mass of Sirius B was known to be about equal to that of the Sun, so its density had to be extremely large. The high density of Sirius B was confirmed in 1925, when the gravitational redshift of its spectral lines was measured. This measurement also provided early observational support to Einstein's general theory of relativity. White dwarfs occur both as single stars and in binary systems. Their spectral lines are broadened by the strong gravitational field at the surface. In some white dwarfs the spectral lines are further broadened by rapid rotation. Strong magnetic fields have also been observed.

White dwarfs have no internal sources of energy, but further gravitational contraction is prevented by the pressure of the degenerate electron gas. Radiating away the remaining heat, white dwarfs will slowly cool, changing in colour from white to red and finally to black.

The cooling time is comparable to the age of the Universe, and even the oldest white dwarfs should still be observable. Looking for the faintest white dwarfs has been used as a way to set a lower limit on the age of the Universe.

white dwarf A compact, dense star nearing the end of its evolutionary life. A white dwarf is a star that has exhausted the thermonuclear fuel (hydrogen and helium) in its interior, causing nuclear burning to cease. When a star of 1 solar mass or less exhausts its nuclear fuel, it collapses under gravity into a very dense object about the size of Earth. Initially, the surface temperature of the white dwarf is high, typically 10,000 K or more. However, as cooling by thermal radiation continues, the object becomes fainter and fainter until it finally degenerates into a non visible stellar core derelict called a BLACK DWARF.

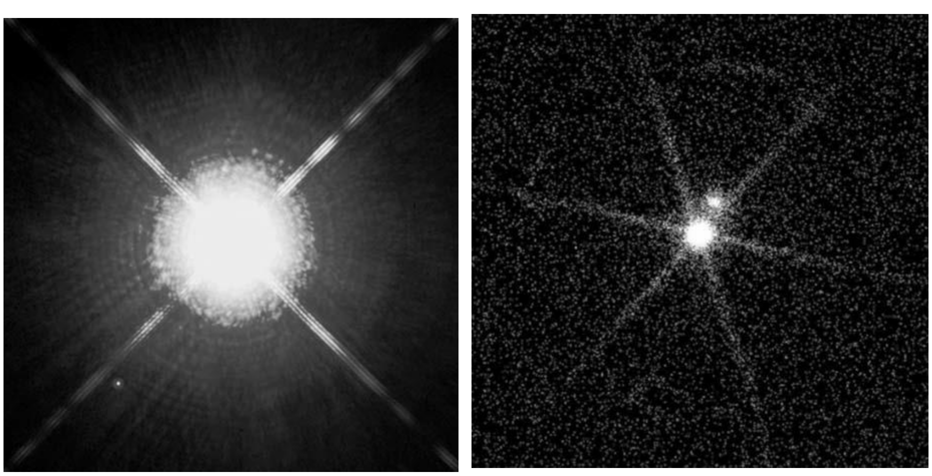
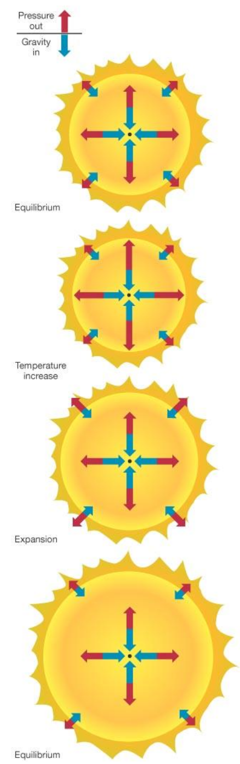


Fig. Two views of the best-known white dwarf Sirius B, the small companion to Sirius. On the left, a picture in visible light by the Hubble Space Telescope. Sirius B is the tiny white dot on lower left from the overexposed image of Sirius.

Stellar Evolution

The set of processes that gradually transform a newly-formed, chemically homogeneous (main sequence) star into sequential phases like red giant or supergiant, horizontal branch star, Cepheid variable, and so forth onward to death as a white dwarf, neutron star, or black hole. The star changes its brightness and surface temperature (and so can be followed on an HR Diagram). It also burns a sequence of nuclear fuels from hydrogen and helium burning (for low mass stars) on up to silicon burning (for stars of more than about 10 solar masses). It is thechanges of chemical composition resulting from these nuclear reactions that are primarily responsible for the changes in luminosity and temperature that we see.

Leaving the Main Sequence

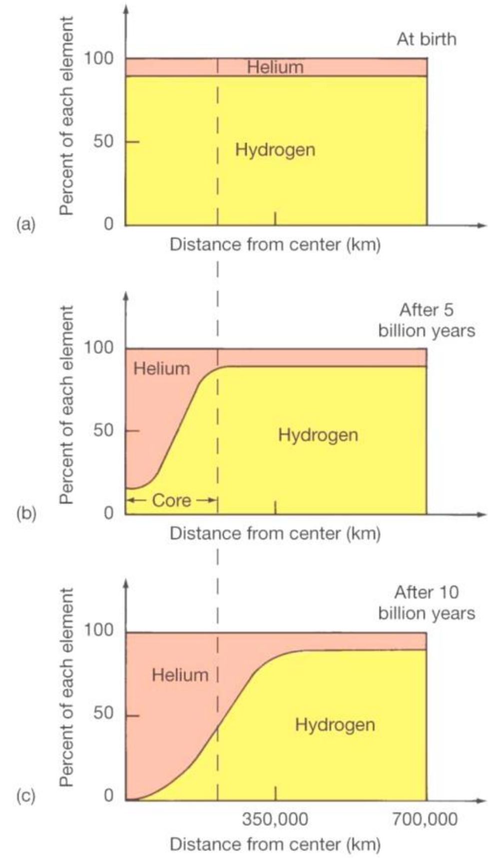


We cannot observe a single star going through its whole life cycle Even short-lived stars live too long for that Observation of stars in star clusters gives us a look at stars in all stages of evolution allows us to construct a complete picture.

During stay on the Main Sequence any fluctuations in a star's condition are quickly restored the star is in Hydrostatic equilibrium.

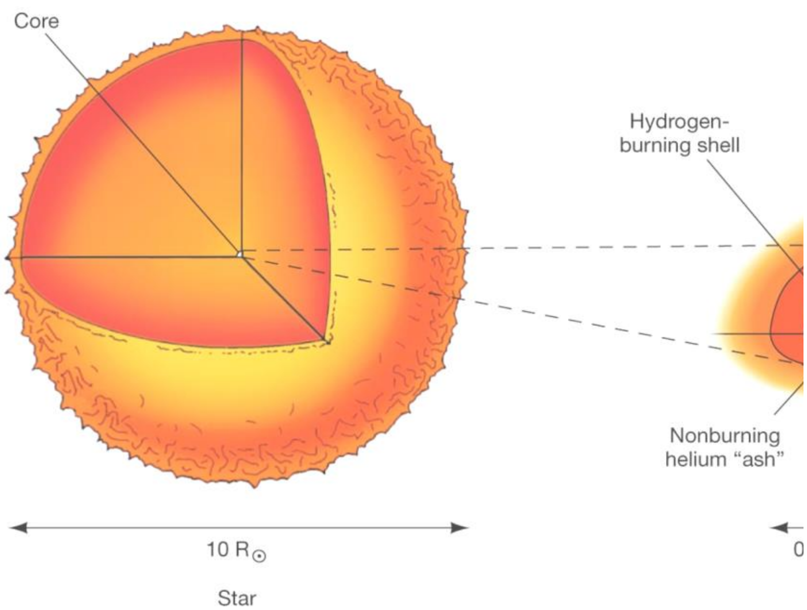
Eventually, as hydrogen in the core is consumed the star begins to leave the Main Sequence Star's evolution from then on depends very much on the mass of the star Low-mass stars go quietly High-mass stars go out with a bang!

Evolution of a Sun-like Star



Even while on the Main Sequence the composition of a star's core is changing amount of helium in the core is increasing

As the fuel in the core is used up, the core contracts the core then begins to collapse Hydrogen begins to fuse outside the core



Stages of a star leaving the Main Sequence

TABLE 20.1 Evolution of a Sun-like Star

Stage	Approximate Time to Next Stage (Yr)	Central Temperature (10^6 K)	Surface Temperature (K)	Central Density (kg/m^3)	Radius (km)	Radius (solar radii)	Object
7	10^{10}	15	6000	10^5	7×10^5	1	Main-sequence star
8	10^8	50	4000	10^7	2×10^6	3	Subgiant branch
9	10^5	100	4000	10^8	7×10^7	100	Helium flash
10	5×10^7	200	5000	10^7	7×10^6	10	Horizontal branch
11	10^4	250	4000	10^8	4×10^8	500	Asymptotic-giant branch
12	10^5	300	100,000	10^{10}	10^4	0.01	Carbon core
		—	3000	10^{-17}	7×10^8	1000	Planetary nebula*
13	—	100	50,000	10^{10}	10^4	0.01	White dwarf
14	—	Close to 0	Close to 0	10^{10}	10^4	0.01	Black dwarf

** Values refer to the envelope.*

Stellar Lifetimes on the Main Sequence

table 21-1 Approximate Main-Sequence Lifetimes

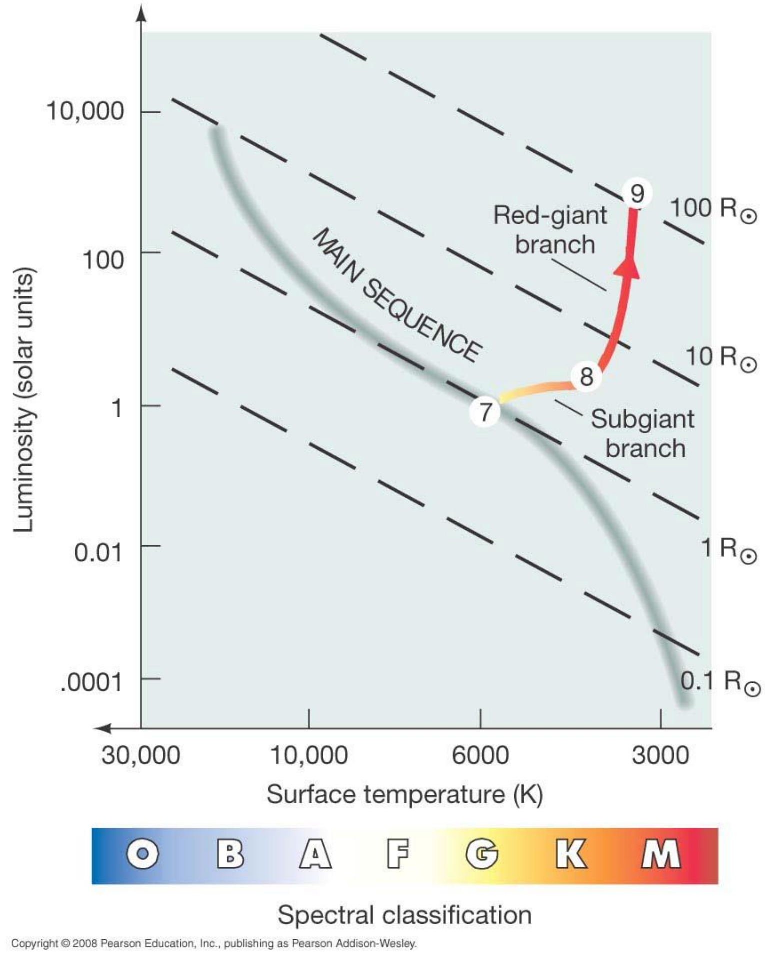
Mass ($M_\odot$)	Surface temperature (K)	Spectral class	Luminosity ($L_\odot$)	Main-sequence lifetim (10^6 years)
25	35,000	O	80,000	4
15	30,000	B	10,000	15
3	11,000	A	60	800
1.5	7000	F	5	4500
1.0	6000	G	1	12,000
0.75	5000	K	0.5	25,000
0.50	4000	M	0.03	700,000

The main-sequence lifetimes were estimated using the relationship $t \propto 1/M^{2.5}$ (see Box 21-2).

Stage 9: The Red-Giant Branch

As the core continues to shrink, the outer layers of the star expand and cool it is now a red giant, extending out as far as the orbit of Mercury despite its cooler temperature, its luminosity increases enormously due to its large size

The red giant stage on the H-R diagram



Stage 10: Helium fusion

Once the core temperature has risen to 100,000,000 K, the helium in the core starts to fuse, through a three-alpha process:

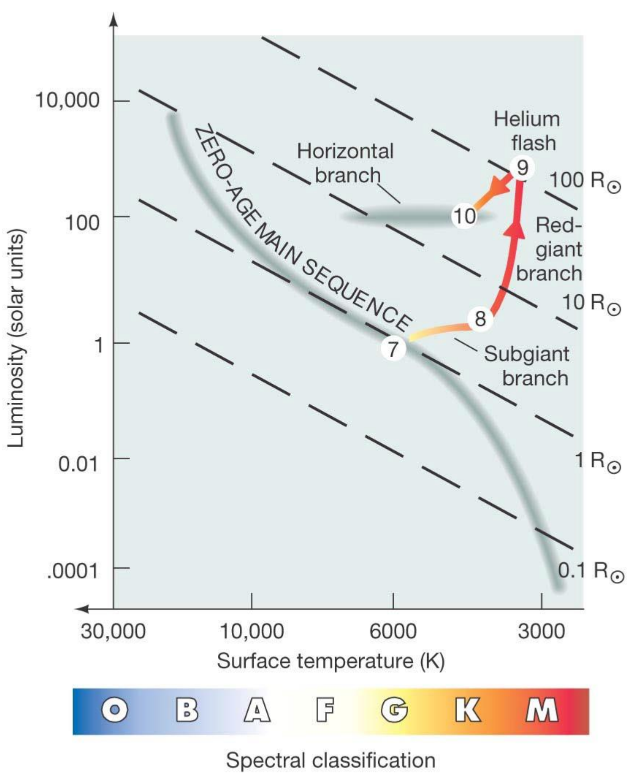
$$^{4}He + {}^{4}He \rightarrow {}^{8}Be + energy$$

$$^{8}Be + {}^{4}He \rightarrow {}^{12}C + energy$$

The ^{8}**Be** nucleus is highly unstable and will decay in about 10^{-12} **s** unless an alpha particle fuses with it first this is why high temperatures and densities are necessary

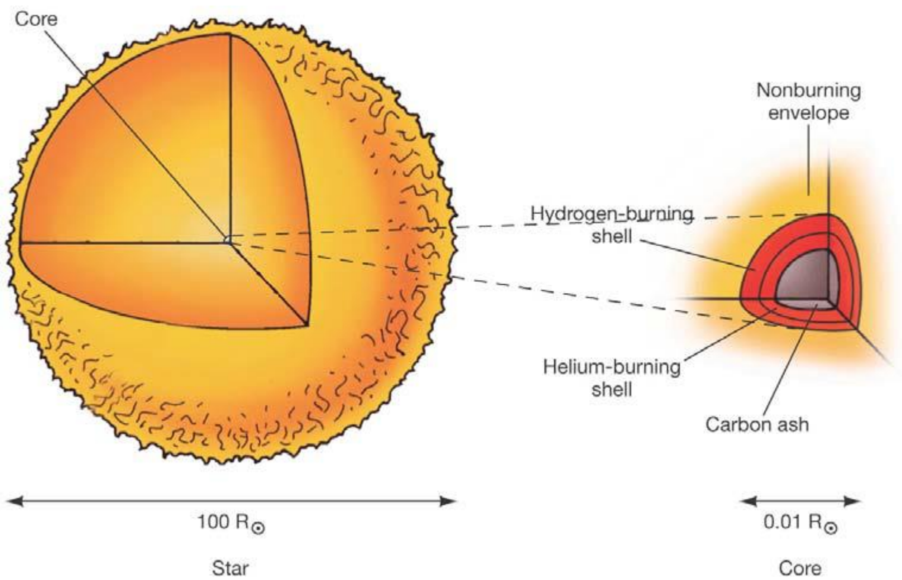
Helium flash

Pressure within the helium core is almost totally due to "electron degeneracy" two electrons cannot be in the same quantum state, so the core cannot contract beyond a certain point this pressure is almost independent of temperature when the helium starts fusing, the pressure cannot adjust helium begins to fuse extremely rapidly within hours the enormous energy output is over, and the star once again reaches equilibrium

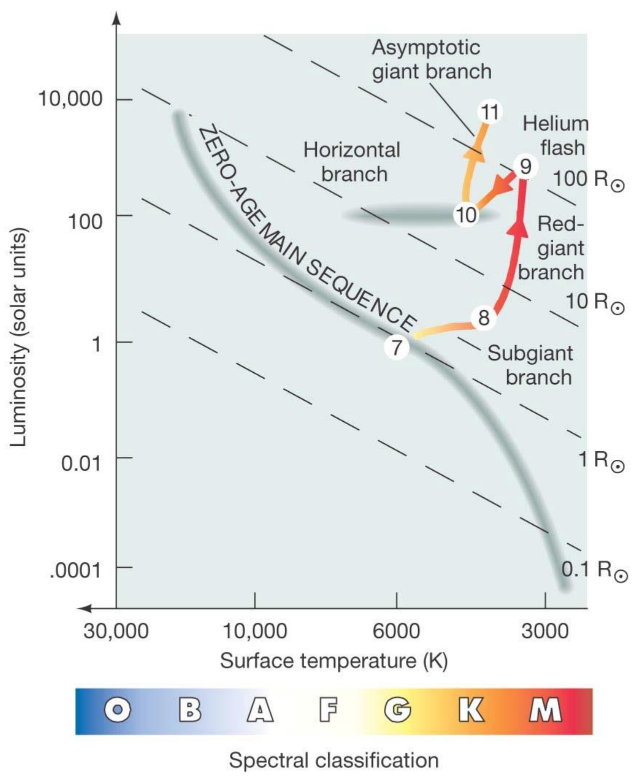


Stage 11: Back to the giant branch

As the helium in the core fuses to carbon, the core becomes hotter and hotter, and the helium burns faster and faster the star is now similar to its condition just as it left the Main Sequence, except now there are two shells



The star has become a red giant for the second time



The CNO Cycle

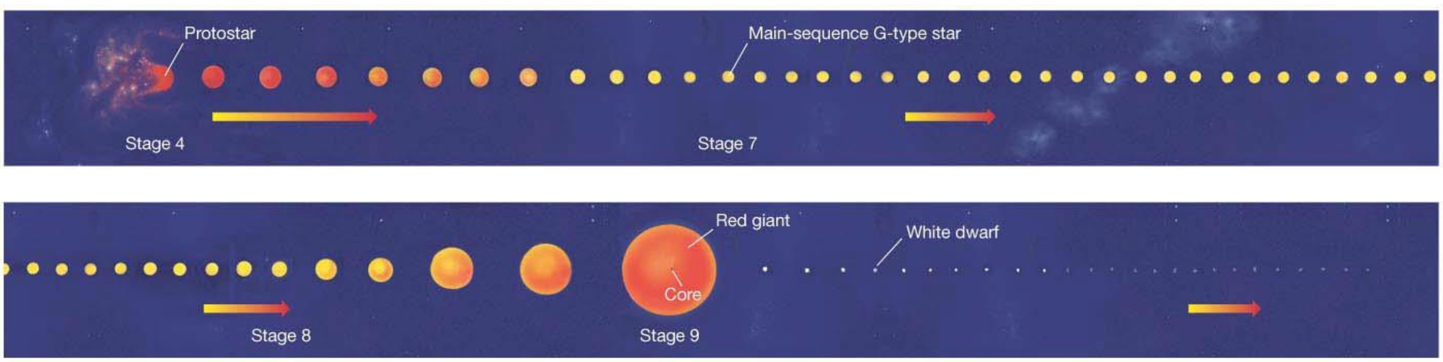
The proton–proton cycle is not the only path stars take to fuse hydrogen to helium

At higher temperatures, the CNO cycle occurs:

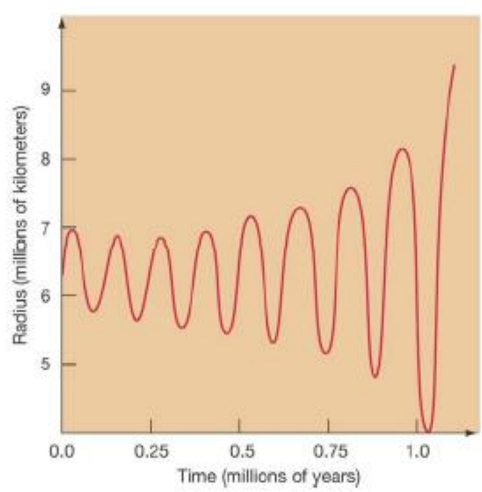
$$
\begin{aligned}
{}^{12}C + {}^{1}H &\rightarrow {}^{13}N + \text{energy} \\
{}^{13}N &\rightarrow {}^{13}C + \text{positron} + \text{neutrino} \\
{}^{13}C + {}^{1}H &\rightarrow {}^{14}N + \text{energy} \\
{}^{14}N + {}^{1}H &\rightarrow {}^{15}O + \text{energy} \\
{}^{15}O &\rightarrow {}^{15}N + \text{positron} + \text{neutrino} \\
{}^{15}N + {}^{1}H &\rightarrow {}^{12}C + {}^{4}He
\end{aligned}
$$

In stars more massive than the Sun, whose core temperatures exceed 20,000,000 K, the CNO process is dominant

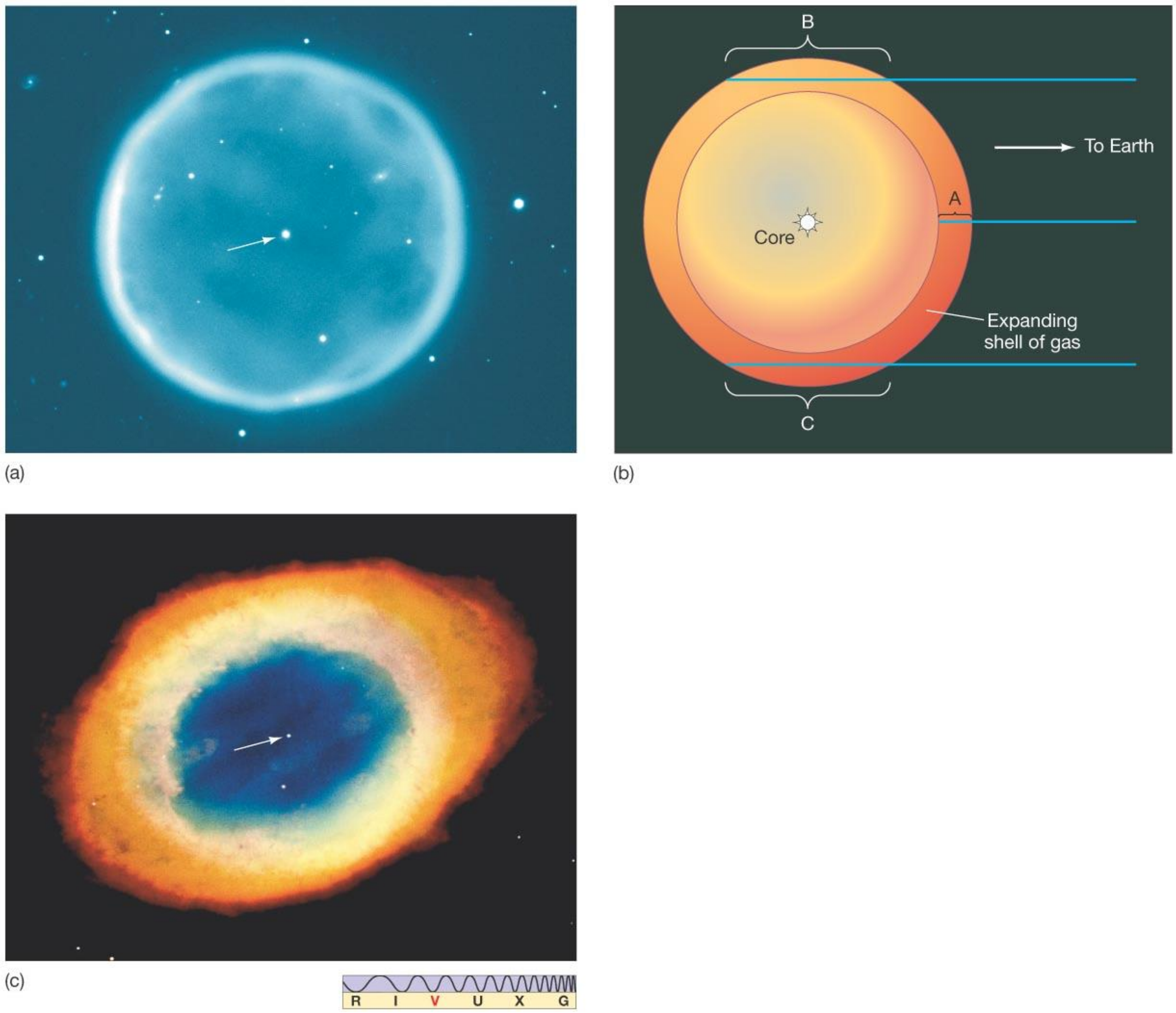
The Death of a Low-Mass Star



There is no more outward fusion pressure being generated in the core, which continues to contract the outer layers become unstable and are eventually ejected



The ejected envelope expands into interstellar space, forming a planetary nebula

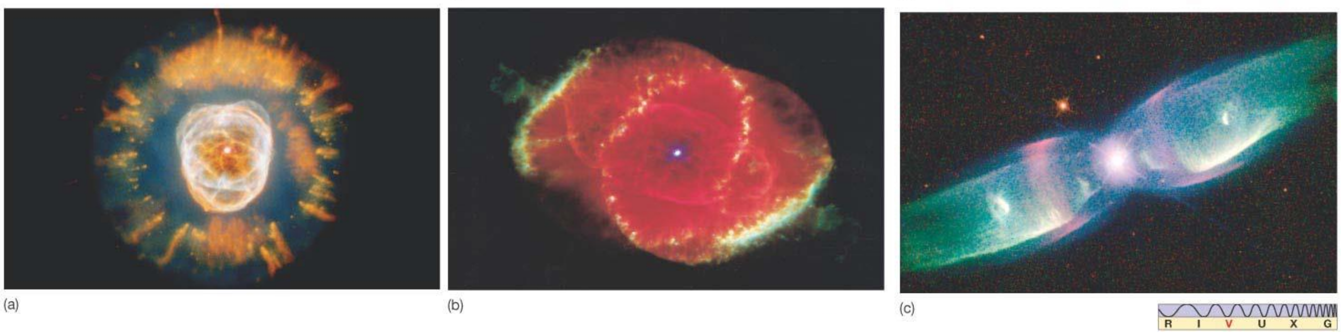


The star now has two parts:

1. A small, extremely dense carbon core

2. An envelope about the size of our solar system – Planetary Nebula (nothing to do with planets)

Planetary nebulae can have many shapes as the dead core of the star cools the nebula continues to expand and dissipates into the surroundings

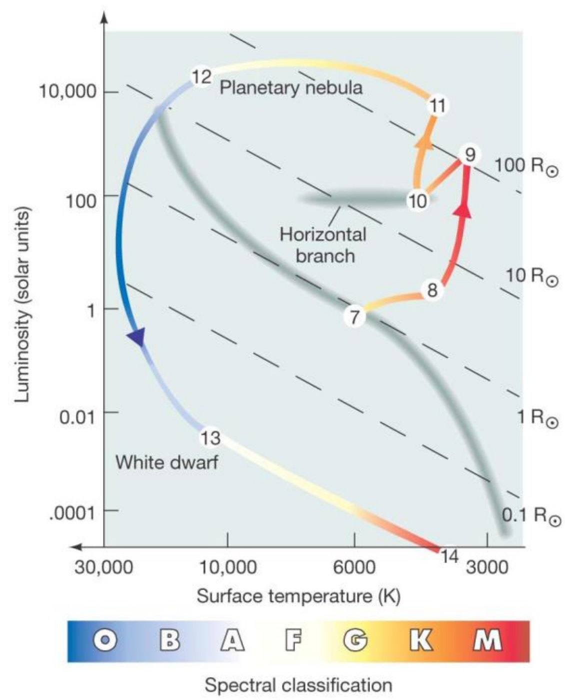


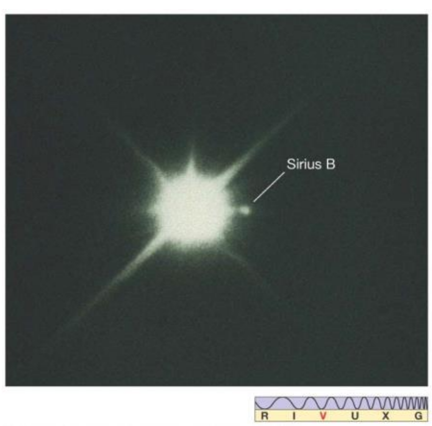
Stages 13 and 14: White and black dwarfs

Once the nebula has gone, the remaining core is extremely dense and extremely hot, but quite small it is luminous only due to its high temperature.

The small star Sirius B is a white- dwarf companion of the much larger and brighter Sirius A

As the white dwarf cools its size does not change significantly, it simply gets dimmer and dimmer, and finally ceases to glow





3.8 STELLAR PARALLAX (TRIGONOMETRIC) AND THE UNITS OF STELLAR DISTANCES

Stellar parallax is a measure of the star's distance. The diameter of the Earth's orbit around the Sun is used as a basic line of such measurements and the corresponding parallax is called the trigonometric parallax. Thus, the trigonometric parallax of a star may be defined as the angle p (in seconds of arc) subtended at the star σ by the mean radius 'a' of the Earth's orbit round the Sun. This is also called heliocentric or annual parallax. A nearby star σ is photographed from the position E of the Earth against the background of the far off stars. Similar observations are also made after six months, from the position E' of the Earth on the other end of the base line. From these observations, the angle subtended by the diameter EE' (or base line) at the star σ is measured, comparing the change in the position of σ during the six months with respect to the fainter far off stars whose shift in position due to the motion of the Earth is negligible. One-half of this angle gives the measure of annual parallax p of the star. The sets of observations at E and E' are repeated when the Earth comes back to this position after a year and the average of a number of measurements is taken as the parallax of the star. As the stars are too far away, the parallax of a star is a very small angle and its value has been found to be always less than 1″ of arc even for the nearest stars.

From Fig. 3.1, we have

$$p_{rad} = \frac{a}{d}, \text{ where } d = \text{distance of the star.}$$

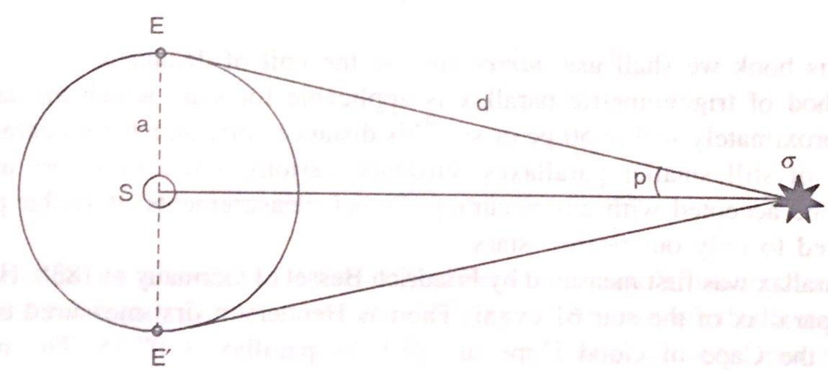


FIGURE 3.1 The annual parallax of a star.

But

$$1_{rad} = 206{,}265'' \text{ of arc.}$$

Therefore,

$$\frac{p''}{206{,}265} = \frac{a}{d},$$

since the parallax of a star is expressed in seconds of arc, or

$$d = \frac{206{,}265\, a}{p''} \tag{3.12}$$

The radius 'a' of the Earth's orbit is called the *astronomical unit* (a.u.) and its value is approximately equal to 150×10^6 km. This is the simplest unit of astronomical distances. If in Eq. (3.12) we take the numerator, i.e., 206,265 a as the unit of distance, then in this unit

$$d = \frac{1}{p''} \tag{3.13}$$

This unit of distance is called a *parsec* (*pc*) and its value in cgs unit is about 3×10^{18} cm. If now in Eq. (3.13) we put $p = 1$, we get $d = 1$. *Thus, a parsec is a distance at which the radius of the Earth's orbit (1 a.u.) subtends an angle of one second of arc.* In other words, parsec is a distance at which the star must be situated in order to exhibit 1 second of parallax, using 1 a.u. as base line. The relation (Eq. 3.13) further shows that the stellar parallaxes are inversely proportional to their distances. Thus, we have

$$1 \text{ parsec} = 206{,}265 \text{ a.u.} \approx 3.26 \text{ light years} \approx 3 \times 10^{18} \text{ cm} \tag{3.14}$$

(To have a proper idea of this distance, the reader may note that if he starts in a vehicle moving at the rate of 1 km s^{-1}, he will take one million years to travel one parsec.)

The relation between a parsec and a light year shows that if distances are measured in light years, Eq, (3.13) will change to

$$d = \frac{3.26}{p''} \tag{3.15}$$

Throughout this book we shall use *parsec* (pc) as the unit of distances.

The method of trigonometric parallax is applicable for stars which are astronomically nearer, i.e. approximately within 50 pc or so. This distance corresponds to a parallax of 0″.02. Measurements of still smaller parallaxes introduce various, kinds of errors and thus such results cannot be accepted with any accuracy. Direct measurements of stellar parallaxes are therefore limited to only our nearest stars.

Stellar parallax was first measured by Friedrich Bessel of Germany in 1838. He successfully measured the parallax of the star 61 cygni. Thomas Henderson first measured the parallax of α Centauri at the Cape of Good Hope in 1839. Its parallax is 0″.75. The nearest star is Proxima Centauri and has got the largest known parallax of 0″.785 with a distance of about 1.3 pc (or 4.24 light year). The parallax of the brightest star, Sirius, is about 0″.379 and its distance is about 2.67 pc (or 8.7 light year).

3.10 STELLAR MOTIONS

Apparently stars seem to be fixed in space. But this is not really the case as they move in space with respect to one another as well as with respect to the Sun. Since the distances of stars are much larger as compared to those of the planets, the stars seem to be stationary with respect to the motion of the planets. Actually the stars are in motion through space with almost equal velocities as those of the planets.

As the stars are at great distances, their relative change of position in the sky is quite small in an interval of one year. The change is noticeable only after a lapse of several years. For this reason stellar motions are observed at an interval of 20 to 50 years and this is measured with respect to the motion of the Sun. The motion of a star through space may be considered to be composed of two components, one along the line of sight of the observer and the other perpendicular to the line of sight. The first component, called the *radial motion,* V_r, determines the speed of recession or approach of the star with respect to the Sun. This is determined by measuring the 'Doppler shift' of the lines in the stellar spectra by the formula $\frac{\Delta\lambda}{\lambda} = \frac{V_r}{c}$, where $\Delta\lambda$ represents the amount of Doppler shift in the spectral line at wavelength λ and c the velocity of light. As the radial velocity of a star is the relative motion between the star and the Sun, necessary corrections for the Earth's motion must be applied. This consists of the corrections due to both the orbital as well as the rotational motion of the Earth. The radial velocity is taken as positive if the star recedes from the Sun, and it is negative when the star approaches it. The velocity is usually expressed in kilometers per second.

The component of the space motion of a star, which is perpendicular to the line of sight is called the *transverse velocity component* or the *tangential velocity component,* V_r. The space velocity, V of a star which is the resultant of the radial velocity (V_r) and the tangential velocity (V_t) is represented in Fig. 3.3.

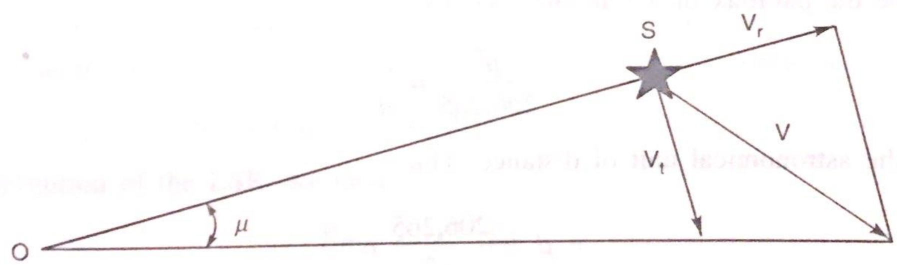


FIGURE 3.3 The diagram showing the different components of the motion of star. The observer is at O and the star is at S.

Thus, V represents the total velocity of a star in kilometers per second and is given by

$$V^2 = V_r^2 + V_t^2 \tag{3.16}$$

The transverse velocity V_t represents the star's movement across the line of sight, i.e. across the direction of V_r. As a result of such a motion, the component V_t describes an angle with the direction of total velocity V. If μ be the angle described at the observer O in one year, then μ is defined as the *proper motion* of the star and is measured in seconds of arc per year. This motion describes the angular rate of displacement in the position as well as in the direction of the star in the sky. Like the radial and the tangential velocities, the proper motion of a star is also considered as observed from the Sun and hence the effect due to Earth's motion must be taken into account. Proper motions of stars are in general very small and range from zero to a few seconds of arc. The largest known proper motion is that of Barnard's star (a star of tenth magnitude) amounting to 10″.25 per year. Large proper motions indicate the high velocity and relative nearness of the star with respect to the Sun so that the parallax can be conveniently measured. Proper motions of the stars in general, are larger than their parallaxes. The mean for all stars which are visible to the naked-eye amounts to about 0″.1.

To calculate the space velocity or total velocity V of a star, we must first know the radial velocity V_r and the transverse velocity V_t. V_r can be calculated from the Doppler formula by measuring the Doppler shift of lines. To calculate V_t, we require the proper motion μ and the parallax p. The former may be known from the value of $N\mu$ taken over a period of N years. The latter can be calculated by applying a suitable method for determining the parallax of a star.

If V_t be the transverse velocity in kms^{-1} of a star at a distance d km and n be the number of seconds in a year, then from Fig. 3.3, we get

$$\frac{\mu''}{206{,}265} = \frac{nV_t}{d}$$

Therefore,

$$V_t = \frac{d\mu''}{206{,}265} \cdot \frac{1}{n} \tag{3.17}$$

If p'' be the parallax of a star, then we also have

$$\frac{p''}{206{,}265} = \frac{a}{d}$$

where a is the astronomical unit of distance. Thus

$$d = \frac{206{,}265}{p''} a$$

Therefore,

$$V_t = \frac{206{,}265}{p''} a \frac{\mu''}{206{,}265} \frac{1}{n} = \frac{\mu''}{p''} \frac{a}{n} \tag{3.18}$$

Using $a \simeq 1.49 \times 10^8$ km, and $n \simeq 3.16 \times 10^7$ seconds, we get

$$V_t = 4.74 \frac{\mu''}{p''} \text{ km s}^{-1} \tag{3.19}$$

From Eq. (3.19) the transverse velocity V_t can be calculated, if the parallax p'' and proper motion μ'' of a star are known. It may be noted that the radial velocity V_r depends only on the shift $\Delta\lambda$ of a line of natural wavelength λ and it is independent of the distance of the star.

Note: In order to determine the proper motion, a star is observed and photographed at an interval of about two decades or more. The photographs are taken against the background of more remote stars or galaxies which remain practically stationary over the observed time periods in the region of the location of the star. From a comparative study of the star images on the two photographic plates, the change in positions of the star during the observed period can be calculated. These observations can however be made only for those stars which are comparatively nearer.

EXERCISES

1. Define parsec; how is it related to astronomical unit (a.u.).
2. The parallax of our nearest star Proxima Centauri is 0″.785. Find its distance in parsecs, light years, astronomical units, miles and kilometres.
3. Calculate the time taken by light to travel to the Earth from the following stars:
   (a) α Centauri with p = 0″.75;
   (b) Proxima Centauri with p = 0″.785;
   (c) 61 Cygni with p = 0″.3;
   (d) Vega with p = 0″.123.

The Doppler shift

It is an everyday experience that the pitch of sound depends on the speed of the object emitting it. The effect was studied by Doppler in the early 19th century and is named after him. In the case of an object which is moving relative to a fixed observer, the sound waves are either lengthened or shortened, depending on whether the object is receding or approaching the observer: the effect of the change of wavelength gives a change in the sensation of the pitch of the sound. It was suggested that the Doppler effect should also occur in connection with light waves but this could only be verified much later in the laboratory when improved optical equipment and techniques became available. It is of historical interest to recall that Doppler erroneously attributed the colours of stars to their motions relative to the Earth. The derivation of the Doppler law here is sufficient for cases where the velocities involved are small relative to the velocity of light.

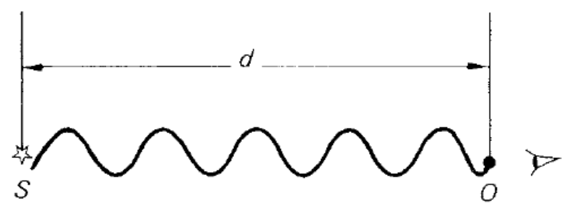


Figure 15.9. A source S at a distance d from the observer O.

Suppose that light is being emitted by atoms in a source, S, with a frequency, ν, and that the waves travel to an observer, O, at a distance, d, from the source (see figure 15.9). After 'switching-on' the source, the first wave arrives at O after a time d/c. During this time the source has emitted ν d/c waves and the apparent wavelength is obviously

$$\lambda = \frac{\text{Distance occupied}}{\text{Number of waves}} = \frac{d}{\nu d/c} = \frac{c}{\nu}.$$

Suppose that the source has velocity, V, away from the observer. During the same time interval as before (d/c), the source moved a distance equal to V × d/c. Thus, the waves emitted during the same interval now occupy a distance equal to

$$d + \frac{Vd}{c}.$$

The apparent wavelength, λ', of the radiation is, therefore, given by

$$\begin{aligned} \lambda' &= \frac{d + Vd/c}{\nu d/c} \\ &= \frac{c}{\nu} + \frac{V}{\nu} \\ &= \lambda + \frac{V\lambda}{c}. \end{aligned} \tag{15.24}$$

Therefore,

$$\lambda' - \lambda = \Delta\lambda = \frac{V\lambda}{c}$$

so giving

$$z = \frac{\Delta\lambda}{\lambda} = \frac{V}{c}. \tag{15.25}$$

The difference, $\Delta\lambda$, between the observed wavelength and the wavelength that would have been observed from a stationary source is called the **Doppler shift**. It is *positive* ($\lambda' > \lambda$) for an object which is receding from the observer and *negative* for an object which is approaching the observer. Thus, lines present in a spectrum of the light from a moving source are shifted towards the *red* for a *receding* object and are shifted towards the *blue* for an *approaching* object.

Equation (15.25) is valid only for $V \ll c$. For velocities near that of light, such as for some galaxies and quasars, the concepts of relativity theory need to be applied giving a Doppler shift which is nonlinear in V. Under this circumstance, the apparent wavelength is written as

$$\lambda' = \gamma\lambda\left[1 + \frac{V}{c}\right] \tag{15.26}$$

where γ is the Lorentz factor:

$$\gamma = \left[1 - \left(\frac{V}{c}\right)^2\right]^{-\frac{1}{2}}.$$

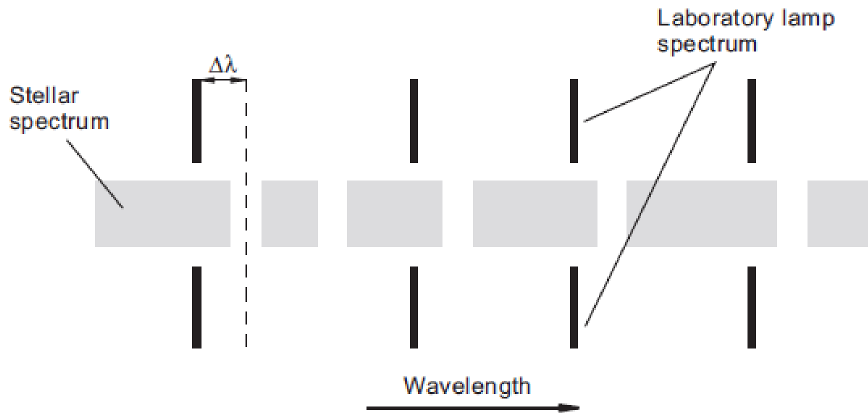


Figure 15.10. A schematic diagram displaying a stellar Doppler shift. Absorption lines within the stellar continuum are matched with displaced emission line spectra of a laboratory reference lamp imposed above and below. Note that the stellar spectrum is redshifted; note also that not all of the stellar lines have matching but displaced lamp lines.

Thus, the relativistic Doppler formula may be written as

$$z = \frac{\Delta\lambda}{\lambda} = \sqrt{\frac{1 + V/c}{1 - V/c}} - 1. \tag{15.27}$$

By comparing the wavelengths of spectral features in the light from celestial objects with standard laboratory wavelengths given by spectral lamps, the Doppler shifts can be measured and velocities deduced. The same optical Doppler effects occur, however, if the source is stationary and the observer moves. Doppler shifts, therefore, represent the **relative motion** between the source and the observer. What is more, the shifts represent only the components of velocity along the line joining the object to the observer. This component is known as the **radial velocity** component and so it can be said that the Doppler shift provides a means of measuring **relative radial velocities**. A schematic spectrum of a star exhibiting the effect of a Doppler shift is illustrated in figure 15.10.

Reference: Astronomy by Roy

Stellar Velocities

The total velocity of a star includes some motion along our line of sight,—that is, either towards or away from us (called the radial velocity)—and some motion across the sky, perpendicular to the radial velocity. This second component is called the proper motion, and it is actually the more difficult measurement to make.

Because optical light with a short wavelength is blue, and long wavelength light is red, when the wavelength of light gets shortened by the Doppler effect, we refer to the change in the wavelength as a Blueshift. When the wavelength of light gets lengthened by the Doppler shift, we refer to the change as a Redshift.

Two additional notes:

It doesn't matter if the source is moving or the observer is moving. That is, if the source of the waves is stationary, but you are approaching it, you will see a blueshift.

The change in the wavelength is proportional to the apparent velocity of the source. That is, the faster the source is moving, the more of a shift you will see.

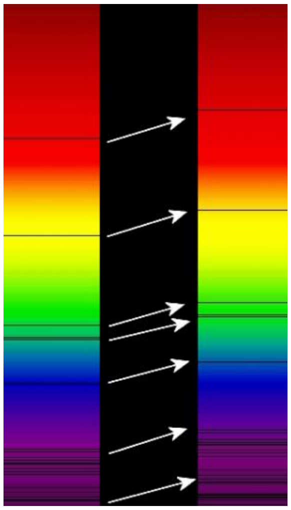


Fig. 4.8: The spectrum from a stationary source on the left, and on the right, the spectrum from the same source with its absorption lines redshifted because of the Doppler Effect.

In practice, astronomers compare the wavelength of absorption lines in the spectrum of a star to the wavelength measured for the same lines produced in the laboratory (for example, the Balmer series lines of hydrogen). The following formula is then used to derive the radial velocity of the star:

$$\Delta \lambda / \lambda_0 = v_r / c$$

In this equation, $\Delta\lambda$ is the difference between the measured wavelength of the line in the star's spectrum and its wavelength in the lab. The rest wavelength is λ_0, which is the wavelength of the spectral line as measured in the lab. The radial velocity of the star is v_r, and c is the speed of light.

For Example, the rest wavelength of the first Balmer line of hydrogen (usually referred to as Hydrogen– α, or just Hα) is 656.3 nanometers . If we measure the Hα in a star to have a wavelength of 657.0 nm, then its radial velocity is:

$\Delta\lambda / \lambda_0$ = (657.0 – 656.3) / 656.3 = 0.001

v_r / c = 0.001

v_r = c × 0.001 = 3.0 x 105 km /sec x 0.001 = 300 km /sec

Since the sign of the velocity is positive, this means that the object is moving at 300 km/sec away from the observer.

This is a very common technique used to measure the radial component of the velocity of distant astronomical objects.

The steps are to

- take the object's spectrum,
- measure the wavelengths of several of the absorption lines in its spectrum, and
- use the Doppler shift formula above to calculate its velocity.

Note that this requires you to know the rest wavelength of the line as measured in the laboratory, which means you need to be able to identify the line in the spectrum of the object even if it has been shifted far from its rest wavelength. This is one of the difficult tasks of observational astronomy.