A short monograph on Stellar Astrophysics

Antonello Piemonte
Marina Rejkuba

Contents

Chapter 1

Some facts about photometry and radiation

In this chapter we will introduce basic definitions about stars, their magnitudes and luminosities, the electromagnetic spectrum and the main radiation mechanisms.

1.1 What are stars?

The stars are huge gas spheres, hundred thousands or millions of times more massive than the Earth. A star such as the Sun can go on shining steadily for thousands of millions of years. This is shown by studies of the prehistory of the Earth, which indicate that the energy radiated by the Sun has not changed by much during the last four thousand million years. The equilibrium of a star must remain stable for such periods.

Before entering such a topic we must review some basic concepts on photometry and spectroscopy. They are needed for a proper understanding of deeper issues like stellar structure and evolution.

1.2 Photometric concepts and magnitudes

The **flux density** gives the energy that crosses a unit area in a unit time; therefore its dimension is $erg\, cm^{-2} sec^{-1}$. The total flux or **luminosity**, is the flux passing through a close surface embracing the source and its dimension is $erg\, sec^{-1}$. If the source radiates isotropically, its radiation at distance r is distributed evenly on a spherical surface whose area is $4\pi r^2$. The relation between Flux and Luminosity is given by $L = 4\pi r^2 F$ where F is the flux density passing through this surface. The physical difference is that, when we are outside the source where radiation is not created or destroyed, the luminosity does not depend on distance. The Flux density on other hand, falls off proportional to $1/r^2$ (fig. 1). For extended objects, like nebulae and galaxies, we use the **Surface Brightness** $B = F/w$ which is the flux density per unit solid angle. B is independent from source distance (fig. 2).

The **magnitude** can be defined in terms of the observed flux density F. We decide that the magnitude 0 corresponds to some preselected flux density F_0. All other magnitudes are then defined by the equation

$$m = -2.5 \log \frac{F}{F_0}$$

In the same way we can show that the magnitudes m_1 and m_2 of two stars and the corresponding flux densities F_1 and F_2 are related by

$$m_1 - m_2 = -2.5 \log \frac{F_1}{F_2}$$

Brighter stars have lower magnitudes values: for example the magnitude of the brightest star, Sirius, is in fact negative, -1.5. The magnitude of the Sun is -26.8 while the faintest star visible with the naked eye have magnitude of +6.

The **apparent magnitude**, which we have just defined, depends on the instrument we use to measure it. We can have different **Magnitude Systems** because different instruments detect different wavelength ranges with different sensitivities. Few important examples:

- *Visual magnitude*: defined by our eyes which is most sensitive to radiation with wavelength around 550 nm

- *Photographic magnitude*: defined by photographic plates which are more sensitive at blue and violet wavelengths

- *Bolometric wavelength*: we would get it if we were able to measure the radiation at all wavelengths. This is very difficult to obtain because of the atmosphere absorption and because we use different detectors at different wavelengths. In practice it is derived from $m_{bol} = m_v - BC$ where BC is the Bolometric Correction. It is zero for solar type stars, and positive for other stars

- *UBV Johnson and Morgan system*: three magnitudes defined by photoelectric photometer plus three filters, U=ultraviolet, B=blue and V=visual (fig. 3). This is a multicolor (multi-filter) photometric system so we can define **colour indices** as the difference between two (different) magnitudes. For example, Vega, which defines the zero point for this system, has $V = 0.03$, $B - V = U - B = 0.00$.

Thus far we have discussed about apparent magnitudes, but they do not tell anything about the true brightness of a star. The quantity measuring the intrinsic brightness of a star is the **absolute magnitude**: it is defined as the apparent magnitude that a star would have at a distance of 10 parsecs. The relation between the apparent magnitude m and absolute magnitude M is given by $m - M = 5 \log r - 5$ where r is the distance expressed in parsecs.

There is an additional factor to consider which is the **extinction**: the space between stars in not completely empty but contains interstellar medium which causes radiation losses. To account for this we rewrite last formula $m - M = 5 \log r - 5 - A$ where $A > 0$ is the extinction (in magnitude) in a certain passband, due to the intervening medium. The other important effect of extinction is the modification of the color indices: if $(B - V)_0$ is the intrinsic colour of the star and $(B - V)$ is the observed one, the we define the colour excess as $E(B - V) = (B - V) - (B - V)_0$ which quantify the variation of the intrinsic colours due to the extinction.

The extinction due to the Earth's atmosphere is characterized by the extinction coefficient k. It depends on the quantities such as the air content of water vapor, aerosol and dust, height of the observatory, etc. and thus changes from place to place. In places that are considered to be good for astronomy k does not vary much with time. For zenithal distances (zdist) $\lesssim 70°$ the linear approximation is valid: $m = m_0 + kz$, where m_0 is the magnitude of a star measured outside the atmosphere and $z = 1/cos(\text{zdist}) = sec(\text{zdist})$.

1.3 Radiation mechanism

We will briefly review the basic concepts related to absorption and emission of radiation. This is fundamental to understand the physical properties of stars. Electromagnetic radiation is emitted or absorbed when an atom or a molecule moves from one energy level to another. An energy level of an atom usually refers to an energy level of its electrons. We know that the energy E of an electron is quantized: an atom can emit or absorb only at certain frequencies ν_{if} corresponding to energy differences between some initial and final states i and f:

$$|E_i - E_f| = h\nu_{if}$$

This gives rise to the **line spectrum** specific for each element. Hot gas under low pressure produces an **emission spectrum**. If the same gas is cooled down and observed against a source of white light (which has a continuous spectrum) the same lines are seen as dark **absorption lines** (see fig. 4).

At lower temperature most atoms are in their lowest **energy state**, the ground state. Higher energy levels are excitation states: a transition from lower to higher state is called excitation. The frequency of the emitted photon is given by $\Delta E = h\nu$. Figure 5 shows different kinds of transitions between energy levels. Figure 6 shows the same for the Hydrogen atom which is the most common element in the universe.

Under particular circumstances we can observe a **continuous spectrum**. This happens in particular for ionizations, recombinations and free-free transitions. The reason for this is that, in those cases, at least one energy state has not quantized energy. Another example is when the pressure of hot gas is increased: in this case atoms bump into each other more frequently and the close neighbors disturb the energy levels. The lines begin to broaden and overlap, thus the spectrum of hot gas at high pressure is continuous. Electric fields also broaden spectral lines (*Stark effect*).

The last important kind of radiation mechanism that we will introduce is the **blackbody radiation**. A black body is an object that does not reflect or scatter radiation shining upon it, but absorbs and re-emits the radiation completely. Such a kind of ideal radiator does not exists in the real world but many objects (including stars) behave very much as if they were blackbodies.

The radiation of a blackbody depends only on its temperature, being independent of its shape, material and internal constitution. The wavelength distribution of the radiation follows the **Planck's law**: the intensity at a wavelength λ of a blackbody at temperature T is

$$B_\lambda(T) = \frac{2hc^2}{\lambda^5}\frac{1}{\exp(hc/\lambda kT) - 1}$$

where h(Planck constant)$= 6.63 \times 10^{-34} Js$, c(speed of light)$= 3 \times 10^8 ms^{-1}$, and k(Boltzmann constant)$= 6.63 \times 10^{-34} JK^{-1}$. The dimension of B_ν is $Wm^{-2}Hz^{-1}sterad^{-1}$. Fig. 7 shows the intensity distribution at different temperatures.

Blackbody radiation can be produced in a closed cavity whose walls absorb all radiation incident upon them (and coming from inside the cavity). The walls and the radiation in the cavity are in equilibrium: both are at the same temperature, and the walls emit all the energy they receive. Since radiation energy is constantly transformed into thermal energy of atoms of the walls and back to radiation, the blackbody radiation is also called **thermal radiation**.

The reason why blackbody radiation is important is that the radiation field deep inside stars behaves like a blackbody! Star's spectra can usually be approximated by blackbody spectra, after the effect of spectral lines has been eliminated.

If we integrate B_ν over all the wavelengths we find that the flux intensity is $F = \sigma T^4$, which is known as the **Stefan-Boltzmann law**, where $\sigma = 5.67 \times 10^{-8} W m^{-2} K^{-4}$ is the Stefan-Boltzmann constant. Now since for a star of radius R and luminosity L we have $L = 4\pi R^2 F$, if the star radiates like a blackbody we can substitute the Stefan-Boltzmann law and we have $L = 4\pi\sigma R^2 T^4$, a very useful formula defining the **effective temperature** of a star: it is the temperature of a blackbody which radiates with the same total flux density as the star.

We can also find the wavelength λ_{max} corresponding to the maximum intensity by differentiating Planck's function with respect to λ and finding zero of the derivative. The result is the **Wien's displacement law**: $\lambda_{max} T = b = 0.0028978$°Km where b is the Wien displacement constant. This formula says that the higher is the temperature the bluer is λ_{max}.

Chapter 2

Properties of Stars

Here we will review the basic properties of stars like their spectra, motions, distances, masses and radii.

2.1 Stellar Spectra

The spectra of stars consist of a continuous spectrum or continuum with narrow spectral lines superposed (fig. 8). The lines in stellar spectra are mostly dark absorption lines, but in some objects, bright emission lines also occur. In a simplified way, the continuous spectrum can be thought of as coming from the hot surface of the star. Atoms in the atmosphere above the surface absorb certain characteristic wavelengths of this radiation, leaving dark gaps at the corresponding points in the spectrum.

Stellar spectra are very important because all our information about the physical properties come more or less directly from studies of their spectra. By studying the strength of various absorption lines stellar masses, temperatures and composition can be deduced.

The spectral classification scheme is based on the temperature of the star (**Harvard Classification** scheme). For different temperatures we have different lines from different elements. Figure 9 shows the variations of some typical absorption lines in the different spectral classes. The spectral sequence is as follow (each class having subclasses denoted by the numbers 0...9, e.g. ...,B2,B1,B0,A9,A8,...):

- Type O: Blue stars, surface temperature 20000 - 35000 K. Spectrum with lines from multiply ionized atoms, e.g. He II, C III, N III, O III, Si V. He I visible, H I lines weak.
- Type B: Blue-white stars, surface temperature about 15000 K. He II lines disappeared, He I starting to appear together with Ca II K line. H I getting stronger. O II, Si II and Mg II lines visible.
- Type A: White stars, surface temperature about 9000 K. The H I lines are very strong and dominate the whole spectrum. Ca II H and K lines getting stronger. He I no longer visible. Neutral metal lines begin to appear.
- Type F: Yellow-white stars, surface temperature about 7000 K. H I getting weaker, Ca II H and K getting stronger. Many other metal lines, e.g. Fe I, Fe II, Cr II, Ti II, clear and getting stronger.

- Type G: Yellow stars like the Sun (which has a spectral type of G2), surface temperature about 5500 K. The H I lines still getting weaker, Ca II H and K lines very strong. Metal lines getting stronger.
- Type K: Orange-yellow stars, surface temperature about 4000 K. Spectrum dominated by metal lines. H I lines insignificant. Strong Ca I and Ca II lines. TiO bands become visible.
- Type M: Red stars, surface temperature about 3000 K. TiO bands getting stronger. Ca I lines very strong. Many neutral metal lines.
- Type C: Carbon stars. Very red stars, surface temperature about 3000 K. Strong molecular bands, e.g. C, CN, and CH. No TiO bands. Line spectrum like the K and M types.
- Type S: Red low (about 3000 K) temperature stars. Very clear ZrO bands. Also other molecular bands.

Note that type C and S represent spectral types parallel to type M, differing in their surface chemical composition.

To summarize, the early (hot) spectral classes are characterized by the lines of ionized atoms, whereas the late (cool) spectral types have lines from neutral atoms (fig. 11). The absorption bands of molecules also appear in the late spectral types. These differences are essentially due to temperature differences.

The Harvard classification only takes into account the effect of the temperature on the spectrum. To be more precise one has to take into account also the luminosity because two stars with the same effective temperature may have very different luminosities. Six **luminosity classes** are defined:

- Ia most luminous supergiants;
- Ib less luminous supergiants;
- II luminous giants;
- III normal giants;
- IV subgiants;
- V main sequence stars (dwarfs)

The luminosity class is determined from spectral lines that depend strongly on the stellar surface gravity, which is closely related to the luminosity. The masses of of giants and dwarfs are roughly similar, but the radii of giants are, by definition, much larger. Therefore the gravitational acceleration $g = GM/R^2$ at the surface of a giant is much smaller than for a dwarf. In consequence the gas density and pressure in the atmosphere of a giant is much smaller. This gives rise to **luminosity effects** in stellar spectrum, which can be used to distinguish between stars of different luminosities. One example is the H I lines: as density increases a fluctuating electric field leads to shifts in the hydrogen energy levels (Stark effect), appearing as a broadening of the lines. Thus the hydrogen lines are narrow in bright stars, and become broader in main sequence stars (fig. 8) and even more so in white dwarfs.

2.2 The Hertzsprung - Russel Diagram

The Hertzsprung-Russel Diagram (HR diagram) illustrates the relation between the absolute magnitudes and the spectral types of stars.

In view of the fact that stellar radii, luminosities and temperatures vary widely, one might have expected the stars uniformly distributed in the HR diagram. In fact it is found that most stars are located along a roughly diagonal curve called the **main sequence** (fig. 12). There are also other features in this diagram that we can easily explain in terms of stellar evolution. For now it is enough to say that most of the stars are concentrated in the main sequence and in the giant branch.

It is important to take into account selection effects: bright stars are more likely to be included in the sample, since they can be discovered at greater distances. If we plot the HR diagram for the stars within a certain distance from the sun we will find no bright main sequence stars and no giants.

2.3 Motions

Many stars seem to move slowly in a direction that does not change with time. This effect is caused by the relative motion of the Sun and the stars through the space; it is called the **proper motion**.

The velocity of a star can be divided in two components (fig. 14), one of which is directed toward the line of sight (the radial component or that radial velocity), and the other perpendicular to it (the tangential component). The tangential velocity results in the proper motion, which can be measured by taking images at intervals of several years or decades. The fastest moving star is the Barnard Star which has a speed of 10.3 arcseconds per year (needs less than 200 years to travel the diameter of a full moon). The radial velocity is measured through the **Doppler effect** which is the change of wavelength of radiation due to the radial motion of the radiation source. A blueshift means that the star is approaching, while a redshift indicates that is receding (fig. 10).

Some useful formulae:

$$V_r = c\frac{\Delta\lambda}{\lambda_0} \qquad V_t = 4.74\mu r$$

(km/s) where r is in parsec and μ in arcseconds per year.

2.4 Distances

The distance of a star is probably the most important physical parameter. The distance unit is the **light-year** which is the distance light travels in one year, or $9.5 \times 10^{15} m$. There are several methods for determining the star's distance. We will explain the most common ones:

- **Trigonometric Parallaxes**: is based on the apparent yearly back and fort movement of stars in the sky, caused by the orbital motion of the Earth. During the course of one year, a star will appear to describe a circle if it is at the pole of celestial sphere, a segment of line if it is in th ecliptic, or an ellipse otherwise. The semimajor axis of the ellipse is called the parallax of the star. It equals the angle subtended by the radius of the Earth's orbit as seen from the stars (fig. 13). So that we can define another distance unit: the parsec. At a distance of one parsec, the semimajor axis of the ellipse is one arc second. One parsec is 3.26 light years. If the parallax p is given in arc seconds the distance (in parsec) is simply $D = 1/p$.

- **Statistical parallaxes**: the velocity of the Sun respect to neighboring stars is about 20 km/s. This means that in one year the Sun moves about 4 AU with respect to the stars. As shown in fig. 16, when the Sun has moved the distance s toward the apex, the direction to the star S appears changed by the angle u. Therefore the distance r is given by $r = ssin\theta/sinu \approx ssin\theta/u$ because the angle u is small. Statistical parallaxes are also called **secular parallaxes**.

- **Photometric Parallaxes**: if we can know the absolute magnitude M of a star, and measure its apparent magnitude m, then we know the distance modulus m-M and the distance. There are several ways to get the absolute magnitude of a star; one of these is to get M by establishing the giant or dwarf nature of the star from its spectrum, enabling it to be placed in the HR diagram. This particular method of photometric parallax is called **spectroscopic parallax**. Another method is scaling an observed HR diagram to fit a theoretical one (MS fitting). This last one is used for determining the distance of star clusters.

- **Kinematic Parallaxes**: it is based on the fact that stars in a cluster all have the same average space velocity with respect to the Sun. Therefore they all appear to be directed to the same point (fig. 17). If ϕ is the angular distance of a given star from the convergent point, V_r is the radial velocity and μ is the proper motion, then the distance can be calculated from

$$d(pc) = \frac{Vsin\phi}{\mu} = \frac{v_r tan\phi}{4.74\mu}$$

  This method is used for young open clusters or associations because they have considerable space velocity.

Fig. 15 compares the reach of different methods used to measure astronomical distances.

2.5 Binary stars and stellar masses

We call binary star systems where two or more stars are gravitationally linked so that they are orbiting around each other. The binary stars are the only known stars with accurately known masses. The masses for other stars are estimated from the mass-luminosity relation (see below). Binaries are classified on the basis of the method of their discovery:

- **Visual binaries**: can be seen as two separate components. The relative position of the components changes over the years as they move in their orbits (fig. 18). In this way the relative orbit of the secondary (the fainter one) can be determined. But the observations give us only the the projection of the orbital ellipse on the plane of the sky (fig. 21). However the shape and position of the true orbit can be calculated because the primary should be located at a focal point of the relative orbit. The deviation of the projected position of the primary from the focus of the projected orbit allows one to determine the orientation of the true orbit. The absolute size of the orbit can be known if we know the distance. Then the total mass of the system can be calculated from Kepler's third law: $M_1 + M_2 \sim a^3/(T^2$ where M is the mass, a is the apparent angular size of the semimajor axis and T is the period. The masses of the individual components can be determined by observing the motions of both components relative to center of mass (fig. 19). If a_1 and a_2 are the semimajor axes of the orbital ellipse for the primary and the secondary, then according to the definition of center of mass $a_1/a_2 = m_2/m_1$ where m_1 and m_2 are the component masses and $a = a_1 + a_2$.

- **Astrometric binaries**: only the orbit of the brighter component about the center of mass can be seen (fig. 20) If the mass of the visible component can be estimated, e.g. from its luminosity, the mass of the invisible companion can also be estimated.
- **Spectroscopic binaries**: appear as single stars in even the most powerful telescopes, but the spectra show a regular variation (fig 22). Either two sets of spectral lines are seen or the Doppler shift of the observed lines varies periodically indicating an invisible companion. If i is the angle between the line of sight and the normal of the orbital plane we have two cases: if one component is so faint that its spectral lines cannot be seen, neither the masses of the components nor the total mass can be determined. The only quantity that we can derive is the so called *mass function* given by $m_2^3 sin^3 i/(m_1+m_2)^2$. If the spectral lines of both components can be observed, it can be demonstrated that we can derive $m_1 sin^3 i$ and $m_2 sin^3 i$. That means we cannot know the actual masses without knowing the inclination. The same holds for the size of the binary orbit (the major semiaxis).
- **Photometric binaries**: in these systems the components of the pair regularly pass in front of each other, causing a change in the total apparent magnitude. Therefore they are also called eclipsing variables. The inclination of the orbit of an eclipsing binary must be very close to 90 degrees. These are the only spectroscopic binaries where the inclination is known and thus where the masses can be uniquely determined. The variation of the magnitude of the eclipsing variables as a function of time is called the light curve. According to the shape of the light curve, they are grouped into three main types: $Algol$, $\beta Lyrae$ and $WUrsae\,Majoris$ type (fig. 23);

¿From the mass derived by binary stars observations it has been found that the mass of main sequence stars become larger, the higher on the main sequence a star is located. One thus obtain an empirical **mass-luminosity relation** (fig. 24) which can be used to estimate stellar masses on the basis of the spectral type. For large masses (above three solar masses), the luminosity is roughly proportional to the cube of the mass. For small masses the dependence is $L \simeq M^{2.5}$. The smallest observed stellar masses are about 1/20 of a solar mass, corresponding to stars in the lower part of the HR diagram. The masses of white dwarfs are less than one solar mass. The mass of the most massive main sequence and supergiant stars are between 10 and 50 solar masses.

2.6 Stellar radii

Direct interferometric measurements of stellar angular diameters have been made for only a few dozen of stars. When the distances are known, these immediately yield the value of the radius. In eclipsing binaries the radius can also be directly measured because the duration of the minima in the light curve depends on the ratio of the stellar radii to the size of the orbit (fig 25). If v is the relative velocity of the two stars, r_s is the radius of the smaller star and r_l is the radius of the bigger one, then $r_s = v/2(t_b - t_a)$ and $r_l = v/2(t_c - t_a)$. Altogether, close to a hundred stellar radii are known from direct measurements. In other cases, the radius must be obtained from the absolute luminosity and effective temperature since $L = 4\pi\sigma R^2 T^4$. Substituting this relation in

$$M_{bol} - M_{bol,\odot} = -2.5 \log \frac{L}{L_\odot}$$

we can express the luminosities in terms of radii and temperatures

$$M_{bol} - M_{bol,\odot} = -5 \log \frac{R}{R_\odot} - 10 \log \frac{T}{T_\odot}$$

If we fix the radius we yield a linear relation between the bolometric magnitudes and $\log T$. So if we use a version of the HR diagram with $\log T$ on the horizontal axis and M_{bol} or $\log L$ on the vertical axis, the lines of constant radius in the HR diagram are straight. They correspond to various values of the radii as shown in fig. 26. The smallest stars are the white dwarfs with radii about 1% of the solar radius, whereas the largest supergiants have radii several thousands times larger than the Sun.

Since the stellar radii vary so widely, so do the densities of stars. The density of giant stars may be one million times less than that of the Sun, whereas the density of a white dwarf is about one million times more.

Chapter 3

Stellar Structure and Evolution

In this chapter we are going to review the structure of the stars, the physical conditions inside them and the source of their energy. Then we will follow their life cycle and what can a star become during its lifetime.

3.1 Stellar structure

As we said before, stars are huge gas spheres which remain in equilibrium for extremely large periods of time. We have four equilibrium conditions governing respectively the distribution of mass, gas pressure, energy production and energy transport in the stars. Each of this condition gives rise to a differential equation, so we have a system with four differential equations which describes the whole status of a star in equilibrium.

The first is the **hydrostatic equilibrium** condition, given by the balance between two forces: the force of gravity, which pulls the stellar material towards the centre, and the pressure force, due to the internal motion of the gas, which pushes the matter outwards. This is expressed mathematically by

$$\frac{dP}{dr} = -\frac{GM_r P}{r^2}$$

The second condition gives the **mass continuity** equation

$$\frac{dM_r}{dr} = 4\pi r^2 \rho$$

it determines the mass contained within a given radius.

The third equilibrium condition expresses the **conservation of energy**, in the requirement that any energy produced in the star has to be carried to the surface and radiated away. If L_r is the energy flux, i.e. the amount of energy passing through the surface r per unit time, and ε is the energy production coefficient, i.e. the amount of energy released in the star per unit time and mass, then

$$\frac{dL_r}{dr} = 4\pi r^2 \rho \varepsilon$$

The rate at which the energy is produced depends on the distance to the centre. Essentially all the energy radiated by the star is produced in the hot and dense core. In the outer layers the energy flux is almost constant.

The fourth equilibrium equation is related to **energy transportation** and its equation gives the temperature gradient $\frac{dT}{dr}$ (the temperature as a function of the radius). The form of this equation depends on the way energy is transported: by *conduction, convection* or *radiation.*

In the interior of normal stars, **conduction** is very inefficient, since the electrons carrying the energy can only travel a short distance without colliding with other particles. Conduction only becomes important in compact stars, white dwarfs and neutron stars, where the free path of electrons becomes very large. In normal stars, conductive energy transport can be neglected.

In **radiative energy transport**, photons emitted in hotter parts of the stars are absorbed in cooler regions, which they heat. The star is said to be in *radiative equilibrium*, when the energy released in the stellar interior is carried outwards entirely by radiation. The radiative temperature gradient is related to the energy flux L_r according to

$$\frac{dT}{dr} = \left(-\frac{3}{4ac}\right)\left(\frac{\kappa\rho}{T^3}\right)\left(\frac{L_r}{4\pi r^2}\right)$$

where a is the radiation constant, c the speed of light, ρ the density, and κ the mass absorption coefficient which gives the amount of absorption per unit mass. It is also called *opacity* because describes how difficult it is for radiation to propagate through the gas. The opacity depends on the temperature, density and chemical composition of the gas. The inverse of the opacity represents the *mean free path* of radiation in the gas, i.e. the distance it can propagate without being scattered or absorbed through various absorption processes (bound-bound, bound-free, free-free). In general the opacity has the functional form of $\kappa = \kappa_0\rho/T^{3.5}$.

The derivative $\frac{dT}{dr}$ is negative, since the the temperature increases inwards. Clearly there must be a temperature gradient, if energy is to be transported by radiation: otherwise the radiation field would be the same in all directions and the net flux would vanish.

If the radiative transfer becomes inefficient, the absolute value of the radiative temperature gradient becomes very large. In that case motions are set up in the gas, which carry the energy outwards more efficiently than the radiation. In these convective motions, hot gas rises outwards into cooler layers, where it loses its energy and sinks again. The raising and sinking gas elements also mix the stellar material, and the composition of the convective parts of a star becomes homogeneous.

The expression for the convective temperature gradient, also called **adiabatic temperature gradient**, is

$$\frac{dT}{dr} = \left(1 - \frac{1}{\gamma}\right)\frac{T}{P}\frac{dP}{dr}$$

where P is the gas pressure and the adiabatic exponent $\gamma = C_p/C_v$ is the ratio of the specific heats at constant pressure and volume, respectively, and can be calculated if the temperature, density and composition of the gas are known.

The convective motions set in when the radiative temperature gradient becomes larger, in absolute value, then the adiabatic gradient, i.e. if either the radiative gradient becomes steep, or if the convective gradient becomes shallow. A steep radiative gradient is expected, if either the energy flux density or the absorption mass becomes large. The convective gradient becomes small if the adiabatic exponent approaches 1.

In order to obtain a well posed problem, four boundary conditions have to be prescribed for the four preceding differential equations. First, there are no energy sources at the centre of mass, thus $M = L = 0$ when $r = 0$. Furthermore, the total mass within the radius R is fixed, $M_R = M$, which defines the value of the radius for a given mass. Finally the temperature and pressure at the stellar surface are very small compared to those in the interior and thus it is usually sufficient to take $T_R = P_R = 0$.

The solution of the basic differential equations give the mass, temperature, density and energy flux as function of the radius. The properties of a stellar equilibrium model are essentially determined once the mass and the chemical composition have been given. This result is known as the **Vogt-Russel theorem** which states that *the mass composition and age of an isolated star uniquely determines its luminosity and effective surface temperature.* Note that, observationally, this is valid only for main sequence stars.

3.2 Stellar energy sources

The source of stellar energy are the **thermonuclear fusion reactions**. In fusion reactions lights elements are transformed into heavier ones. The final products have a smaller total mass then the initial nuclei. This mass difference is released as energy, according to Einstein's equation $E = mc^2$.

In stars with masses similar to the Sun or smaller the energy is produced by the **proton proton (pp) chain**. It consists in the fusion of four protons into one He nucleus, $p^+ + p^+ \rightarrow^4_2 He + p^+ + p^+$, and its production coefficient is $\varepsilon_{pp} \propto \rho T^4$. One chain gives about 26 Mev. It interesting to note that the expected time for a proton to collide with another one is 10^{10} years; It is only thanks to this slowness that the Sun is still burning. If it were faster, it would have burnt out long ago.

At temperatures above 20 million degrees, corresponding to stars with masses above $1.5M_\odot$, the **carbon (CNO) cycle** becomes dominant. This is because its reaction rate increases more rapidly with the temperature $\varepsilon_{CNO} \propto \rho T^{17}$. It also consists in the fusion of four proton into one He nucleus, but it is catalysed by carbon, oxygen and nitrogen. One cycle gives about 25 Mev, slightly less that the pp chain, because more energy is carried away by neutrinos.

Stars burning H at their centers form more than 90% of the observed stars and they are situated in the **Main Sequence**.

As a result of the preceding reactions, the abundance of helium increases. At temperatures above 10^8 degrees the helium can be transformed into carbon in the **triple alpha reaction** and a core of carbon is formed. Once helium burning has been completed, at higher temperatures other reactions become possible, in which heavier elements up to iron are built up. Examples of such reactions are carbon, oxygen and silicon burning. With nuclear fusion it is not possible to form elements heavier then iron because this would require an input of energy. Elements heavier then iron are almost exclusively produced by **neutron capture** during the final violent stages of stellar evolution.

3.3 Stellar evolution

3.3.1 Protostars

Stars are believed to form inside large dense interstellar clouds mostly located in the spiral arms of the Galaxy. Under its own gravity, a cloud begins to contract and fragment into parts that will become protostars. Observations indicate that stars are not formed individually, but in larger groups. Young stars are found in open clusters and in loose associations typically containing a few hundreds stars which must have formed simultaneously.

An interstellar cloud can contract only if its mass is large enough for gravity to over whelm the pressure. The limiting mass for this to happen is the **Jeans mass** given in solar masses by

$M_j \approx 3 \times 10^4 \sqrt{T^3/n}$ where n is the density in atoms per cubic meter. It is thought that star formation begins in clouds of a few thousands solar masses and diameters of about 10 pc.

Interstellar clouds begin to contract because of their passing through a spiral arms which in turn compresses clouds and trigger contraction. Another mechanism might be a nearby supernova explosion. When the cloud contracts, it does not heat up because the liberated energy is carried away by radiation. The contraction and fragmentation continues until the density becomes so large that the energy liberated by contraction can then no longer escape and the temperature will begin to rise. This raises the pressure which in turn stops contraction.

3.3.2 The contraction of stars towards the Main Sequence

Once the contraction of the protostar is stopped the star settles into hydrostatic equilibrium. The gas is almost completely ionized and the further evolution takes place on a time scale which is much longer (millions of years) than the free-fall time scale (hours) valid during the contraction phase. Now the star is located inside a larger gas cloud and will be accreting material from its surroundings. Its mass therefore grows and the central temperature and density increase. The temperature of the star is still low and its opacity correspondingly large. Thus it is convective.

We will describe now the evolution in the HR diagram (fig. 28). Initially the star will be faint and cool and it will reside in the lower far right part of the diagram, outside the figure. During its collapse it will move to the upper right, settling at a point corresponding to its mass on the **Hayashi track**. The Hayashi track gives the location in the HR diagram of completely convective stars. The star will now move downwards, its radius decreases and its luminosity drops. As the temperature increases in its centre, the opacity diminishes and energy begins to be transported by radiation. The mass of the convective region will decrease until finally most of the star is radiative. By then, the central temperature will have become so large that nuclear reactions begin. Previously released energy was potential energy, but now nuclear reactions overtake and the luminosity increases.

The beginning of the **Main Sequence** phase is marked by the start of hydrogen burning in the pp chain.

3.3.3 The Main Sequence Phase

The main sequence (MS) phase is that evolutionary stage in which the energy released by the H burning is the only source of stellar energy. During this stage the star is in stable equilibrium and it structure changes only because its chemical composition is gradually altered by the nuclear reactions. This phase takes the longest part of the life of a star (fig. 31).

Since stars are most likely to be found in the stage of steady H burning, the MS in the HR diagram is richly populated in particular to its low mass end. The more massive upper MS stars are less abundant because of their shorter MS lifetimes.

The stars in the **upper MS** are so massive and their central temperature so high that they can operate the CNO cycle. The energy production is concentrated into the core and the outgoing energy flux is so large that cannot be maintained by radiative transport. Therefore there is a convective core i.e. energy is transported by material motions. These keep the material well mixed and thus the H abundance decrease uniformly with time within the convective region. Outside the core there is radiative equilibrium, i.e. energy is transported by radiation and there are no nuclear reactions (fig. 27b).

The mass of the convective core will gradually diminish as the H is consumed. In the HR diagram the star will slowly shift to the upper right. When the central H supply is exhausted the core of the star will begin to shrink rapidly. The surface temperature will increase and the star will quickly move to the upper left. Because of the contraction of the core, the temperature of the H shell just outside the core will increase enough for the H burning to set in again.

On the **lower MS** the central temperature is lower than for massive stars and the energy is generated by the pp chain. Since the rate of the pp chain is less sensitive to temperature than the CNO cycle, the energy production is spread over a larger region. In consequence the core never becomes convective but remains radiative.

In the outer layers of lower MS stars the opacity is high because of the low temperature. Radiation can then no longer carry all the energy and convection will set in. The structure of lower MS stars is thus opposite to that of the upper MS: the centre is radiative and the envelope is convective (fig. 27b).

As H in the core decreases the star will slowly move upwards in the HR diagram, becoming slightly brighter and hotter, but its radius will not change much. Then the evolutionary track will bend to the right as the H in the core nears its end (fig. 27a). Eventually the core is almost pure He. H will continue to burn in a thick shell around the core.

3.3.4 The giant phase

The MS phase ends when the H is exhausted at the centre. The star then settles in a state in which H is burning in a shell surrounding a He core. The mass of the He core is increased by the H burning in the shell. This leads to an expansion of the envelope of the star which moves almost horizontally to the right in the HR diagram.

In **low mass stars**, as the mass of the core grows, its density will eventually become so high that it becomes degenerate. What does it means? Just before the onset of helium burning the electrons (remember that at these temperatures the atoms are completely ionized, thus most of the matter in the star's core consists of detached nuclei and electrons) are so closely crowded together that any further compression would violate a quantum mechanics law called the **Pauli exclusion principle**. It states that two identical particles cannot simultaneously occupy the same *quantum state*, that oversimplifying means that you can't have two things in the same place at the same time. So the closely packed electrons cannot be squeezed anymore and produce a powerful pressure that resists further core contraction. We say that the helium-rich core of a low-mass red giant is *degenerate* and is supported by **degenerate electron pressure**. This degenerate pressure, unlike the pressure of a perfect gas, does not depend on the temperature.

The whole He core will have now a uniform temperature because of the high conductivity of the degenerate gas. The central temperature will continue to rise until it will be enough for He to burn to C in the triple alpha reaction.

The He burning will further raise the central temperature, but the degenerate core cannot expand, although the temperature increases, and therefore there will be a further acceleration of the nuclear reaction rate. When the temperature increase further the degeneracy of the gas is removed and it will begin to expand violently producing an explosion, the **He flash**. The energy from the He flash is absorbed by the outer layers, and thus does not lead to the complete disruption of the star. After the He flash, the star will settle in a new state where He is steadily burning to C in a nondegenerate core and will find itself on the **horizontal giant branch** in the HR diagram (fig.32).

When the He is exhausted in the core, there are two nuclear burning shells: an inner one where He is burning, an an outer one where H is burning instead.

In **high mass stars** the central temperature is higher and the central density is lower, and the core will therefore not be degenerate. Thus He burning will set in non-catastrophically as the central region contracts. When the central He supply is exhausted, He will continue to burn in a shell.

3.3.5 Final stages

The evolution that follows He burning depends strongly on the stellar mass. The mass determines how high the central temperature can be and the degree of degeneracy when heavier nuclear fuels are ignited.

Stars less massive than $3M_{\odot}$ never become hot enough to ignite carbon burning on the core. At the end of the giant phase, the radiation pressure expels the outer layers which forms a **planetary nebula**. The hot core remains as a **white dwarf**.

In stars with initial masses in the range $3-15M_{\odot}$, either carbon or oxygen is ignited explosively just like helium in low mass stars: there is a **carbon** or **oxygen flash**. This is much more powerful than the helium flash and will make the star explode as a **supernova**, leaving no remnant.

The most massive stars, with masses larger than $15M_{\odot}$, burn all the way to iron ^{56}Fe. All nuclear sources of energy will be then completely exhausted. The star is made up of a nested sequence of shells burning silicon, oxygen, carbon, helium and hydrogen (fig. 29). However this is not a stable state since the cessation of nuclear reactions in the core means that the central pressure will fail and the core will collapse. The energy released in the collapse goes into dissociating the iron nuclei first to helium and then to protons and neutrons. This will further speed up the collapse which will take place in only a fraction of a second and will release immense amount of energy in few seconds as a supernova explosion. If the mass on main sequence was less than about $25M_{\odot}$, the remnant in the inner core will stabilize and become a **neutron star**. If the mass was larger nothing can support the remnant against the pull of the gravity and the final collapse will be complete producing a **black hole**, an object whose gravity is so high that not even light can escape from it.

3.3.6 Planetary Nebulae

As we have seen in connection with stellar evolution, instabilities may develop at the stage of helium burning. In some stars the whole outer atmosphere may be violently ejected into space. A gas shell expanding at 20-30 km/s will be formed around a small and hot (50.000 - 100.000 K) star, the core of the original star.

The planetary nebulae were given their name in the 19th century, because certain small nebulae visually look quite like planets. The apparent diameter of the smallest known planetary nebulae is only a few arcseconds, whereas the largest ones (like the Helix nebula) may be one degree in diameter.

The expanding gas is ionized by ultraviolet radiation from the central star and its spectrum contains bright emission lines due to hydrogen (656.3 nm) and forbidden lines from doubly ionized oxygen (495.9 and 500.7 nm) and ionized nitrogen (654.8 and 658.3 nm). **Forbidden lines** are so called because they are extremely difficult to observe in laboratory. Their transition probabilities are so small that at laboratory densities the ions are de-excited by collision before they had time to radiate. In the diffuse, extremely low-density environment of planetary nebulae, collisions are much less frequent, and thus there is a chance that an excited ion will make the transition to a lower state by emitting a photon.

Planetary nebulae are generally symmetrical in shape and expand rapidly. In few thousands years they disappear in the general interstellar medium and their central star cools to became a white dwarf. The total number of planetary nebulae in our Galaxy has been estimated to be 50000. About 1000 have been actually observed.

Chapter 4

Compact and Variable Stars

In this chapter we will examine more closely some exotic objects left from star's death, and also we will briefly review the numerous family of variable stars.

4.1 Compact stars

The stars in which the density of matter is much larger than in ordinary stars are known as **compact objects**. These include white dwarfs, neutron stars and black holes. In addition to a very high density, the compact objects are characterized by the fact that nuclear reactions have completely ceased in their interiors. Consequently they cannot support themselves against gravity by thermal gas pressure. In the white dwarfs and neutron stars, gravity is resisted by the pressure of a degenerate gas. In black holes, the gravity is completely dominant and compresses the stellar material to infinite density.

4.1.1 White dwarfs

The white dwarfs are quite numerous in space but faint and therefore difficult to find. The best known example (and the first to be discovered, in 1915) is Sirius B, the companion of Sirius.

They occur both as single stars and in binary systems. Their spectral lines are broadened by the strong gravitational field at the surface. In some of them the spectral lines are further broadened by rapid rotation. Depending on their composition they can show different kinds of spectra: solar type, helium rich (depleted in hydrogen), carbon rich (depleted in hydrogen and helium), or they can show no lines at all.

Their typical radius is about 10000 km, but they have masses comparable to that of the Sun. Therefore the central density can be as high as $10^9 kg/m^3$. The mass of a white dwarf cannot be larger than the **Chandrasekar Mass** $M_{Ch} = 1.44 M_{\odot}$. When a massive star reaches the end of its evolution and explodes as a supernova, the simultaneous collapse of its core will not necessarily stop at the density of a white dwarf. If the mass of the collapsing core is larger than the Chandrasekar Mass, the collapse continues to a neutron star.

Since a white dwarf has no internal energy source, it will slowly cool down, changing in color from white to red and finally to black. In the HR diagram they will move to the right, following one of the constant radius lines. There ought be a large number of black dwarfs in the Galaxy.

4.1.2 Neutron Stars

A neutron star is formed in the explosion and collapse of a supernova. It will initially rotate very rapidly (several hundreds times per second) because its angular momentum is unchanged while its radius is much smaller than before.

Neutron stars are difficult to study optically because their luminosity in the visible region is very small, typically $10^{-6}L_{\odot}$. For instance the Vela pulsar has been observed at a visual magnitude of about 25. On the other side in the radio region, they are a very strong pulsating source.

The typical diameters of neutron stars are about 10 km. Unlike ordinary stars they have a well defined solid surface. The atmosphere above it is a few centimeters thick. The upper crust is a metallic solid with a density growing rapidly inwards. Most of the star is a neutron superfluid, and in the centre, where the density exceeds $10^{18}kg/m^3$ there may be a solid nucleus of heavier particles.

The composition of a neutron star is explained by the extreme density of the matter which renders the gas degenerate. Matter, electrons and protons get squeezed so much that they combine to form neutrons. Matter then consists of a neutron porridge, mixed with about 0.5% electrons and protons.

Neutron stars are supported against gravity by the pressure of the degenerate neutron gas, just as white dwarfs are supported by degenerate electron gas pressure. The equation of state is the same except that the electron mass is replaced by the neutron mass.

The theory of neutron stars was developed in the 1930's, but the first observations were not made until the 1960's. At that time a new type of rapidly pulsating radio sources, the **pulsars**, were discovered and identified as neutron stars.

The **pulsars** (PULSAting Radio Source) have the following characteristics: periods between 0.25s and 2s, period accuracy set them as clocks comparable or better than best atomic clocks, their periods are increasing gradually (characteristic lifetime about 470 million years). The basic pulsar model (fig. 34) consists of a rapidly rotating neutron star with a strong dipole magnetic field (two poles) that is inclined to the rotation axis at an angle θ.

The radiation emitted by pulsars is **synchrotron radiation** which is produced when relativistic electrons spiral along magnetic field lines (fig. 33). The radiation is emitted in a narrow radio beam in the instantaneous direction of motion of the electron. As the neutron star rotates, these radio waves sweep through space in a way reminiscent of the light from a rotating lighthouse beacon.

The supernova explanation for the origin of the pulsars is consistent with the fact that only 1% of known pulsars belong to a binary system, while at least half of the stars in the sky are found in multiple stellar systems. Pulsars also move much faster through space than do normal stars, as though the explosion's violence flung them through space while disrupting any nearby companion.

4.1.3 Black Holes

If the core of a collapsing star exceed the **Oppenheimer-Volkov mass** $M_{OV} \approx 2M_{\odot}$ not even the degenerate neutron gas pressure can resist the gravitational force. There is no stable final state: the force of gravity will dominate over all other forces and the star will become a **black hole**. It is black because not even light can escape from it. According to classical mechanics the escape velocity from a body of radius R and mass M is $v = \sqrt{(2GM)/R}$. This is greater then the speed of light when the radius is smaller than a critical radius called the **Schwarzschild radius** $R_S = 2GM/c^2$. Because the mass of a black hole must be larger than M_{OV} the radius of the smallest black hole is about 5-10 km.

The properties of black holes have to be studied on the basis of the **general theory of relativity**. In the theory of relativity each observer carries with him his own local measurement of time. If two observers are at rest with respect to each other at the same point, their clocks go at the same rate. Otherwise their clocks rates are different and the apparent course of events, too.

An **event horizon** is a surface through which no information can be sent out, even in principle. A black hole is surrounded by an event horizon at the Schwarzschild radius. Near the event horizon, the different time definitions become significant. An observer falling into a black hole reaches the centre in a finite time, according to his own clock, and does not notice anything special as he passes through the event horizon. However to a distant observer, he never seems to reach the event horizon; his velocity of fall seems to decrease towards zero as he approaches the horizon.

The slowing down of time also appears as a decrease in the frequency of light signals: the frequency observed by an infinitely distant observer approaches zero for radiation emitted near the event horizon.

The tidal forces become extremely large near a black hole so that any material falling into it will be torn apart. All particles are destroyed near the central point and the final state of matter is unknown to present-day physics.

A black hole can have only three observable properties: mass, angular moment and electric charge. At present, the only known way to directly observe a black hole is by means of the radiation of gas falling into it. For example, if a black hole is part of a binary system, gas streaming from the companion will settle into a disc around the hole. Matter at the inner edge of the disc will fall in the hole. The accreting gas will lose a considerable part of its energy as radiation, which should be observable in the x-ray region. Some very rapidly and irregularly varying x-ray sources are promising black holes candidates.

4.2 Variable Stars

Stars with changing magnitudes are called variable stars. Strictly speaking all stars are variables; as we have seen the structure and brightness of a star change as it evolves, although these changes are usually small.

Variables are usually divided in three main types: *pulsating, eruptive and eclipsing variables.*

In the **pulsating variables** the light variations are due to the expansion and contraction of the outer layers. These variables are giants and supergiants that have reached an unstable stage in their evolution. The main cause of light variation is the periodic variation of the surface temperature. In fact we know that the luminosity of a star depends sensitively on its effective temperature $L \propto T_e^4$. Thus a small change in temperature leads to a large brightness variation. Also the diameter of the star change with luminosity so that the wavelengths of the spectral lines change due to Doppler Effect. The observed velocities are in the range of 40-200 km/s.

Normally stars are always adjusting in a stable hydrostatic equilibrium. If their outer layers expand, the density and temperature decrease. The pressure then becomes smaller and the force of gravity compresses the gas again. Unless energy can be transferred to the gas motions, these oscillations will be damped.

The origin of the pulsation mechanism rely on ionization zones where hydrogen and helium are partially ionized. Therefore the opacity becomes larger when the gas is compressed and if the ionization zones are at a suitable depth in the atmosphere an instability occurs: the energy absorbed during compression and released during expansion of an ionization zone can drive an oscillation. Stars with surface temperatures of 6000-9000 K are subject to this instability and the corresponding section on the HR diagram is called the **instability strip** (fig. 30). In this instability strip we have several families of pulsating variables: *RR Lyrae, Classical Cepheids, Dwarf Cepheids, Mira-like* and more.

The **eruptive variables** are usually (but not always) faint stars ejecting mass. They are mostly members of close binary systems in which mass is transferred from one component to the other. They have no regular pulsation. Instead, sudden outburst occur in which material is ejected into space. The scale of the outbursts goes from small local eruptions (*flare stars*) to the explosion of the entire star (*supernovae*).

The **eclipsing variables** are binary systems in which the components periodically pass in front of each other. In these variables the light variations do not correspond to any physical change in the stars.

Chapter 5

The Milky Way

This chapter will summarize the overall picture of the Galaxy and its components: stars, star clusters and interstellar medium.

5.1 What is the Milky Way?

On a clear moonless night a nebulous band of light can be seen stretching across the sky. This is the **Milky Way**, also called the Galaxy. It is the reflection of a huge system consisting mostly of stars. The stars of the Milky Way form a flattened disk-like system. In the direction of the plane of the disc huge numbers of stars are visible, whereas relatively few are seen in the perpendicular direction. The faint light from distant stars merges into a uniform glow and therefore the Milky Way appears as a nebulous band to the naked eye.

The Milky Way is a **spiral galaxy** surrounded by an extended halo of objects with the Sun located in the disc about 8 kpc away from the centre (fig. 35). The centre appears to be in the constellation of Sagittarius.

In studying the structure of the Milky Way it is convenient to choose a spherical coordinate system so that the fundamental plane is the symmetry plane of the Milky Way (fig. 37).

In order to study the structure of the Milky Way we need to know the distribution in space of its constituents such as stars, star clusters and interstellar matters. So we can understand now why the distance of stars is a primary parameter for such a study. Also the stellar density, determined by means of star counts on photographic plates, is a crucial parameter.

What we know today is that the young objects like OB associations and open clusters, are strongly concentrated in the galactic plane. Old objects, particularly globular clusters, have an almost spherical distribution about the centre of the Milky way.

5.2 Stellar Populations

Studies of the motions of the stars in the Milky Way have revealed that the orbits of stars moving in the galactic plane are almost circular. These stars are usually young and they also contain a relatively large amount of heavy elements, about 2%-4%. Also the interstellar medium shows the same kinematic properties. On the basis of their motions and chemical composition, the interstellar medium and the youngest stars are collectively referred to as **population I**.

Outside the plane of the Milky Way, an almost spherically symmetric halo extends out to 50 kpc and beyond. Here the stars are old, up to 15 Gy, and there is very little interstellar matter.

The stars are also very metal poor and their orbits may be very eccentric. On the basis of these criteria, one defines stars of **population II**. Typical population II objects are the globular clusters and the RR Lyrae stars.

The origin of the **spiral structure** is due to a wavelike variation of the density of the disc. The spiral pattern rotates as a solid body with an angular velocity about half that of the galactic rotation, while the stars and gas in the disc pass through the wave. This explains why young objects are found in the spiral arms. As gas passes through the wave it is strongly compressed. The internal gravity of the gas clouds then becomes more important and causes them to collapse and form stars.

5.3 Star Clusters

Some Stars are collected in clusters and associations. **Associations** are groups of very young stars. They are usually identified on the basis of bright main sequence stars or of T Tauri stars. According to the type one speaks of *OB associations* or of *T Tauri associations*. The first one are very massive stars that stay on the main sequence only few millions years. The latter are younger stars that are in the process of contracting towards the main sequence.

Associations are dispersing rapidly. They have large amounts of interstellar matter and are strongly concentrated in the spiral arms of the Milky Way. Therefore they are very young, population I stars.

Open Clusters contain up to few 100 stars and are also population I objects. Their HR diagrams show very well defined and narrow main sequence; there are only a few giants. Usually one can clearly see the point in the HR diagram where the main sequence ends and bends towards the giant branch. This point depends strongly on the age of the cluster. It can therefore be used to determine the ages of open clusters. The HR diagram can also be used to determine their distances through the method called the main sequence fitting.

Globular Clusters usually contain about 10^5 stars and are more concentrated (fig. 38) and have central densities about 10 times larger than open clusters. Globular clusters are among the oldest objects in the Milky Way which contains about 200 of them. Mostly they are found in the halo, but some of them are associated with the bulge.

The HR diagram of a typical globular cluster is shown in fig. 39. The main sequence only contains faint red stars; there is a prominent giant branch, horizontal branch and asymptotic branches are (usually) clearly seen. The main sequence is lower than that of the open clusters. The metal abundances are much lower in globular clusters.

5.4 Interstellar matter

Although most of the mass is concentrated into stars, interstellar space is not completely empty. It contains gas and dust, in the form both of individual clouds and of a diffuse medium. Interstellar space contains typically about one gas atom per cubic centimeter and 100 dust particles per cubic kilometer.

Altogether about 10% of the mass of the Milky Way consists of interstellar gas. It is concentrated mainly in the galactic plane and the in spiral arms. The dust (a better name would be smoke, since the particle size are much smaller than in terrestrial dust) constitutes about 1% of the gas and is confined in a thin, about 100 pc, layer in the galactic plane.

Dust can be concentrated in individual clouds, which appear like star poor regions or **dark nebulae** against the background of the Milky Way. If a dust cloud is near a bright star it will scatter i.e. reflect the light of the star producing bright **reflection nebulae**. It is composed mainly of water ice and silicate.

The most important observations of the interstellar medium are made at radio and infrared wavelengths, since the peak of the emission often lies at these wavelengths. The upper limit on the total mass of interstellar medium can be derived on the basis of its gravitational effects.

The interstellar dust is responsible for some effects (fig. 36). The first one is **light extinction**, which was already introduced before. It is due to dust grains that have diameters near the wavelength of the light. Such particles scatter light extremely efficiently (gas also cause extinction by scattering but its scattering efficiency per unit mass is much smaller) leading to a reduced intensity in the original direction of the propagation. Extinction by dust can be due also to absorption, where the radiant energy is transformed into heat, which is then reradiated at infrared wavelengths corresponding to the temperature of the dust particles. This temperature is about 10 - 20 K and the corresponding wavelength is $300 - 150 \mu m$. Near a hot star the temperature of the dust can be 100 - 600 K and the maximum emission is then at $30 - 5 \mu m$.

Light extinction varies strongly with direction. Although a value of $A = 2 mag/kpc$ is used as average, visible light from the galactic centre (distance 10 kpc) is dimmed by 30 magnitudes. Therefore the galactic centre cannot be observed at optical wavelengths. On the other side the total extinction towards the galactic poles may be less than 0.1 mag.

Light extinction also varies with the wavelength: it is larger for shorter wavelengths, causing the other effect, **reddening** of the light of stars. It is found that $A \approx 1/\lambda$. For this reason the light from distant stars is redder than would be expected on the basis of their spectral class. The spectral class is defined on the basis of the relative strengths of the spectral lines which are not affected by extinction.

The mass of the interstellar gas is hundred times larger then that of dust. Nevertheless it is less easily observed since the gas does not cause a general extinction of light. It can only be observed on the basis of a small number of spectral lines like the sodium D_1 and D_2 lines in the optical region, or the hydrogen 21 cm line in the radio wavelengths.

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