Diploma in Astronomy, 2002–2003

Foundations of Astronomy

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Figure references in italics are to the course textbook, *Universe*, 6th edition, by Freedman and Kaufmann.

Lecture of 1 October, 2002.

Introduction

In this module, we are looking at the Sun, stars, galaxies and everything else outside the Solar System. We'll consider observations not only in visible light but in other radiation such as the radio and X-rays. We start with how we can label the objects we are studying, to identify them and to know where to point a telescope to observe them.

Constellations (*See Figure 2.2, in three parts*)
A constellation is a pattern that appears in the bright stars when the human mind joins up the dots. The constellations have no real physical meaning. In general, the stars all are at different distances from us so what we see depends on our particular point of view; and different civilisations didn't always see the same pictures in the dots anyway.

Originally a pattern in the stars, a constellation now is defined as an area of the sky; the whole sky is covered with 88 constellations (you don't have to memorise them!) and there are no gaps. Anything in that part of the sky can be said to be in that constellation.

The Naming of Stars

A few hundred of the brightest stars have individual names, such as Sirius, which have come to us from antiquity. Beyond this, Greek letters are used, more or less in order of brightness within a constellation. As the brightest star in Canis Major, Sirius is α CMa (using a standard three-letter abbreviation). We soon run out of letters and there is a numbering system, introduced by Flamsteed, based on position in the constellation. E.g., 61 Cygni is a well-studied nearby star, in the constellation Cygnus.

For most stars now, we use the entry number in some catalogue. The best-used is the HD (Henry Draper) catalogue, with more than 300,000 entries. For the vast number of fainter objects that we have now detected, generally the position in some coordinate system is used.

Other Wavebands.

When the first radio surveys were made around 1950 they detected only a few sources. These were labelled by letters with the constellation name, e.g., Cygnus A is the brightest radio source in the constellation Cygnus. Later, with greater sensitivity many more sources were found and catalogues were prepared. The third Cambridge catalogue labels some hundreds of sources, e.g. 3C273 which is the nearest quasar.

In X-rays, for the first strong sources found again the constellation name was used. Cygnus X-1 is the strongest X-ray source in Cygnus, and was later found to be the same active galaxy that is also the radio source Cygnus A.

Celestial Coordinates

To establish a position on Earth we use two numbers, the latitude and longitude, defined using a great circle, the Equator (perpendicular to the rotation axis) and a fixed point on it, arbitrarily chosen to be given by the meridian through Greenwich.

In the same way, a coordinate system in the sky, fixed among the stars, can be defined from a great circle in the sky and a fixed point on it (*See Fig 2-9 and Box 2-1*). We can used the celestial equator, which is everywhere directly overhead at the Earth's equator, and take as fixed point the place where the Sun crosses the equator each year, from south to north, which is at the time of the Vernal Equinox. This point is called the First Point of Aries, and written as Υ. A wobble in the Earth's motion called precession means that Υ moves very slowly through the stars (it takes 26,000 years to go right round the equator) and it now no longer is in the constellation Aries.

Corresponding to latitude, we have the Declination, which is measured in degrees, 0° to $+90^\circ$ for north and 0° to -90° for south of the equator.
Corresponding to longitude, we have the Right Ascension which is measured round the equator in hours, 0 to 24 hours. This is convenient in practice for as the Earth rotates one hour of A passes overhead in one hour of time. (This is an hour of *sidereal time*, measured against the stars rather than the Sun).

Sirius has coordinates RA = 6h 41m, Dec $=-16^\circ\ 35'$. Note that we have hours, minutes and seconds of time (hms), and degrees, minutes and seconds of arc. $1^\circ = 60'$ (minutes) and $1' = 60''$ (seconds).

The Nature of Light.

In a rainbow, sunlight is converted by the raindrops into a spectrum of different colours. The same thing happens in a prism, and this was seen in an experiment carried out by Newton (*Figures 5.3, 5.4*). A beam of sunlight from a small hole in a shutter passes through a prism and gives the spectrum. Selecting just one colour, with another hole in a screen, and passing it through another prism just that one colour emerges. This proved that the different colours are present in the original sunlight and not created inside the prism. The "white" light is not broken up but spread out into the different colours as the colours merge into each other in a continuous spectrum.

Lecture of 8 October, 2002.

The Nature of Light (continued).
Other experiments over the 200 years from Newton's time in the late 17th century showed that light behaves as a form of wave motion. Waves are familiar in nature, the easiest to picture being surface waves on water. In the case of light, its interactions are related to the electrical and magnetic properties of matter and it is described as an electromagnetic wave.

A wave has a characteristic *wavelength*, the distance between successive peaks; each of the colours in the rainbow has its own wavelength.
We also use the *frequency* of the wave, the number of peaks that pass a fixed point per second. The time for one wave to pass equals the wavelength divided by the speed, distance divided by speed, or *time* = *wavelength*/*speed*. The frequency, the number of waves per second, is one over this, $frequency$ = *speed*/*wavelength*.

We can use symbols for these quantities; the standard ones are greek letters:
λ(*lambda*) = *wavelength*, ν(*nu*) = $frequency$, c = *speed*.
λ is a length, in metres say, written m, c a speed in metres per second , written as m/s or m s^{-1}. The frequency ν is waves per second and is given the name Hertz (Hz).
Thus, ν in Hz = (c in m/s)/(λ in m), or $\nu = c/\lambda$.
Equivalently, $\nu\lambda = c$, $\lambda = c/\nu$.

Powers of 10.
The wavelengths of visible light are much smaller than 1 metre. We can also have lengths that are larger than 1 metre. It is helpful to have labels that help us to cope with these.

We have 1 kilometre (km) = 1000 m = 10 × 10 × 10 m, which we can write as 10^3 m. Also, 1 metre = 1000 millimetres (mm) = 10^3 mm, or 1 mm = 1/1000 m = 10^{-3} m.

Continuing to shorter lengths, 1 micron (μm) = 10^{-6} m (where μ is the greek letter mu), 1 nanometre (nm) = 10^{-9} m.

The Electromagnetic Spectrum
Visible light has wavelengths of order hundreds of nm, from about 400 nm (violet) to 700 nm (red) (*See Figure 5.3*).

The electromagnetic spectrum extends indefinitely in both directions (*See Figure 5.7*):

700 nm to $\sim$ 1mm	Infrared
$\sim$ 1mm to $\sim$ 10 cm	Microwave
$\sim$ 10 cm onwards	Radio
400 nm to $\sim$ 10 nm	Ultraviolet
$\sim$ 10 nm to $\sim$ 10^{-2} nm	X-Rays
$\sim$ 10^{-2} nm to 10^{-6} nm and shorter	γ Rays

Here, $\sim$ means about, and γ is the greek letter gamma.

In the laboratory, X-rays are produced by a particular type of process in atoms and γ rays by certain radioactive atomic nuclei. In Astronomy, they have come to mean *any* radiation with this sort of wavelength, regardless of how it is produced.

In empty space, all electromagnetic radiation travels at the same speed, $c = 3 \times 10^8$ m s^{-1} (approximately). With frequency $\nu = c/\lambda$, for visible light the frequency is a very large number and it is not normally used. Frequency is much used in Radio Astronomy, with frequencies involving

MegaHertz (MHz) = 10^6 Hz	corresponding to	$\lambda = 300$ m
GigaHertz (GHz) = 10^9 Hz		$\lambda = 0.3$ m

Lecture of 15 October, 2002.

The Nature of Light (continued).
Many results show that light behaves like a wave motion. Early in the 20th century, it was found that in certain experiments it behaves as though it were a stream of particles. The important names are Planck and Einstein and you can read about them in the textbook. Light is not either a wave or a particle. It has properties of both and some experiments bring out its wave nature, others its particle nature.

The particles of light are called *photons*. For radiation of frequency ν, the energy carried by each photon is proportional to ν and can be written as $E = h\nu$ where h is called Planck's constant. With frequency in Hz and E expressed in the standard energy unit Joules (J), $h = 6.625 \times 10^{-34}$ J s. This is very small; the effects of the particle nature of light are far too small to be seen in everyday experience.

The Joule itself is inconveniently large to be used as a unit in atomic physics and spectroscopy and instead we often work with a unit called the electron volt (eV), where 1 eV $= 1.602 \times 10^{-19}$ J.
Visible light photons have energies of a few eV. X-ray astronomers work with units of MeV and GeV, that is 10^6 eV and 10^9 eV.

Transmission of the Atmosphere.
Visible light gets through the Earth's atmosphere virtually without loss. There are some wavelength ranges (called *windows*) in the infrared where the radiation comes through, and radio astronomy can also be done from the ground, but for wavelengths shorter than about 300 nm and for most of the infrared we have to observe from above the atmosphere. (*See Figure 6-27*).

The reason is the absorption of the radiation by molecules in the Earth's atmosphere. Moving into the UV, it is a form of oxygen called ozone (O_3) that starts absorbing at the longest wavelength, so the destruction of ozone where it is present high in the atmosphere lets through more of the potentially- harmful UV radiation—hence the concern about the ozone hole in the atmosphere. In the IR and microwave regions, many molecules absorb but water vapour H_2O is the most important one; water is present mainly in the lower atmosphere where weather is formed and high mountain observatories such as Mauna Kea in Hawaii are above much of this IR absorption.

At radio wavelengths longer than about 20 m the radio waves are reflected away by the Earth's ionosphere (a region in the upper atmosphere). Photon energies are very low and it is most unlikely that there is any useful astronomical information at such wavelengths.

Colours of the Stars.
The Sun has a yellowish tinge, and other stars can be red, or white, or blue. This is caused by differences in their surface temperatures.

The detailed form of the spectrum emitted by a star is rather complicated, but to quite a good approximation it can be thought of in terms of the simpler spectrum emitted by an idealised object known as a *black body*.

Black Body Radiation.
A black body is defined to be an object that perfectly absorbs all radiation incident on it, of all wavelengths. It can be represented in the laboratory by the inside of an enclosure with a very small aperture, so that any photon that finds its way in is reflected around inside and is most unlikely to come out again. If the enclosure is maintained at a particular temperature, it emits radiation which has the following properties derived from experiment (*cf. Figures 5-10, 19-7*):

1. The spectrum of black-body radiation depends only on the temperature and is independent of the material of which the enclosure is made.

2. The intensity has a maximum at some particular wavelength, which we may call

λ_{max}, with less intensity at longer or shorter wavelengths.

3. The total amount of energy emitted is proportional to the fourth power of the temperature, $E \propto T^4$. Thus, if the temperature is doubled, 16 times as much energy is given out. This is called Stefan's Law (or the Stefan-Boltzmann Law). For unit area of a surface, it can be written as $E = \sigma T^4$ where σ is the greek letter sigma and is called Stefan's constant.

4. The wavelength λ_{max} of the maximum intensity shifts to shorter wavelengths for higher temperatures, and follows the rule $\lambda_{max} = 0.0029/T$ which is called Wien's Law.

5. Black-body curves for different temperatures never cross.

The temperature scale used in these relations is in degrees Kelvin or degrees absolute. Temperature is a measure of the amount of motion in a system; the faster things move, the higher the temperature. The imagined situation where nothing is moving at all defines the absolute zero of temperature. It is about $-273°$ C and nothing can be colder. This is the zero of the Kelvin scale; a change of one Kelvin degree is the same as one Celsius degree so $0°$ C is $+273$ K.

It turns out that the spectrum of a star is not much different from that of a black body of suitable temperature. For the Sun, the intensity curve is quite close to a black body curve for 5800 K (see Figure 5-11 but note that the solar spectrum shown is not at all accurately drawn).

Lecture of 22 October, 2002.

Stars and Black Body Spectra

Just as for the Sun, to a first approximation we can represent the spectra of other stars by black body curves. The best way to fix the temperature is to do it precisely by measuring the total energy emitted by the star, at all wavelengths, and using Stefan's Law $E = \sigma T^4$ to estimate the temperature. This is called the *effective temperature*, sometimes written T_{eff}. Loosely it is the surface temperature of the star; being a cloud of gas, a star does not have a well-defined surface with a single temperature, but T_{eff} is a reasonable average of the actual gas temperatures found in the atmosphere of the star.

The Range of Astronomical Temperatures.

The Sun has $T_{eff} \sim 5800$ K and looks yellowish. It is a typical middle-of-the-road star. Some other stars are cooler; it becomes hard to observe stars much below about 3000 K as they are very faint and the radiation is in the infrared. Planetary temperatures are typically a few hundred degrees K. The hottest stars have temperatures around 100,000 K. Such objects are very rare, but they are very luminous so they can be seen at great distances.

The biggest range of temperatures is for the different forms of interstellar matter. The densest interstellar clouds are less than 10 K; at the other extreme intergalactic gas, between the galaxies in some clusters of galaxies, is of order 10^8 K.

The Brightness of Stars.

It is evident from a glance at the sky, or in a photograph, that the stars have different brightness. There are two reasons for this—there are differences in intrinsic brightness, and they are at different distances.

To see how the effect of distance works, suppose we have a collecting area of 1 m^2; this could be the effective area of the mirror of a large telescope. We can measure the radiation that falls on this collecting area but not what falls on either side. At distance d from a star, the total area of a surface centred on the star and right round it is $4\pi d^2$ metres^2. Here π (greek pi) is the constant 3.14.... A fraction $1/(4\pi d^2)$ of the radiation from the star hits our detector. At twice the distance, we detect one quarter the amount. Thus, the farther away the star, the less of its light is detected so it appears fainter.

Measuring Distances.
Until recently, surveying on Earth used triangulation. With a baseline of known length we can measure the angles from either end to a distant point and (using trigonometry) work out its distance. We can do the same for the stars. The largest baseline available is the diameter of the Earth's orbit, called two Astronomical Units or 2 AU. By radar ranging of the planets we can set the scale of distances in the Solar System and find the AU in metres. A star is observed many times from around the orbit (to increase the accuracy) and so if it is near enough its distance can be found (*See Figure 19-2*).

This phenomenon is called *parallax* and this word is also used as the name of a quantitative measure. The parallax of a star is defined to be the difference in angle at which it is seen from two points 1 AU apart on a baseline perpendicular to the direction of the star.

These angles are very small. The nearest star, Proxima Centauri, has parallax 0.772 seconds of arc. All other stars have smaller values of the parallax and most stars are too far away to have a measurable parallax angle at all.

Lecture of 29 October, 2002.

Units of Distance.
One *parsec* (pc) is the distance of an object with parallax one arcsec (we don't actually know any object at this distance). An object at 2 pc has parallax 1/2 arcsec, and so on. p (arcsec) $= 1/d$ (parsec).

If we know the AU in km, from trigonometry we can find the value of the parsec. We can compare this with the *light year*, the distance light travels through empty space in one year, at speed 3×10^8 m s^{-1}.

$$\begin{array}{rcl} 1 \text{ AU} &=& 1.496 \times 10^8 \text{ km} \\ 1 \text{ ly} &=& 9.46 \times 10^{12} \text{ km} \;=\; 6.32 \times 10^4 \text{ AU} \\ 1 \text{ pc} &=& 3.09 \times 10^{13} \text{ km} \;=\; 3.26 \text{ ly} \;=\; 2.06 \times 10^5 \text{ AU} \end{array}$$

The distance to Prox Cen is

$$d = \frac{1}{p} = \frac{1}{0.772} = 1.3 \text{ pc} = 4.2 \text{ ly}.$$

Thus, light from this star takes 4.2 years to reach us.

For greater distances we use kpc (10^3 pc) and Mpc (10^6 pc). Distances in our Milky Way galaxy are expressed in kpc—the disk of the Galaxy is about 30 kpc across. Apart from a few nearby ones, other galaxies are at Mpc distances.

Parallax measurements are very difficult because of the small angles involved. A great improvement came in the early 1990s with the European satellite *Hipparcos* which measured parallaxes for more than 100,000 stars to distances of several hundred pc. For more distant stars, and for galaxies, we can't measure distance directly. Indirect methods can be used, based on intrinsic properties of the stars, and you'll meet some of these methods later on.

Distance measurement is a step-by-step process, using one method to calibrate distance to some type of object that can then be used to greater distances. The first two steps are (1) radar ranging of planets to find the AU, and (2) trigonometric parallax measurement for nearby stars.

Stellar magnitudes.
How bright is a star in the sky?

The ancient system classified the stars in six groups by magnitude, with first magnitude for the brightest stars seen and sixth magnitude for those just visible to the naked eye.

They go in equal steps. It happens that, in our physiology, we respond to stimuli such as light and sound in such a way that changes actually by the same factor appear to us as equal steps. Thus, if the brightness of a source of light doubles, and then doubles again, we perceive these as changes by the same amount. So any difference of 1 magnitude corresponds to a difference always by the same factor in the intensity.

It turns out that a typical first magnitude star has about 100 times the intensity of light of a sixth magnitude star. This is a chance result. We now *define* a difference of 5 magnitudes to correspond exactly to a factor 100 in intensity. This is for differences 1st–6th, 2nd–7th, 20th–25th, and so on.

A difference of one magnitude corresponds to a factor 2.512 in the intensity. We can see that this is so, for five such steps gives

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

a factor 100 as it should be.

Taking sixth magnitude to correspond to stars listed as sixth magnitude in the old catalogues based on visual observation, we can use this relation to determine the magnitudes of other stars from intensity measurements. What were originally called 'first magnitude' stars actually have a range of intensity.

A star that is a factor 2.512 brighter than first magnitude is called magnitude 0, one brighter than that by the same factor is magnitude -1, and so on. The brightest star, Sirius, has magnitude -1.4. The plot in *Figure 19.6* shows the range that we have, from the Sun to the faintest stars that are currently observable.

For two stars, magnitudes m_1 and m_2, intensities received I_1 and I_2, then the fundamental relation is

$$\frac{I_2}{I_1} = 10^{0.4(m_1 - m_2)}.$$

Notice how intensity changes by a factor, magnitude by addition. We can see how this formula works for two cases:
If $m_1 - m_2 = 5$ then $I_2/I_1 = 10^{0.4 \times 5} = 10^2 = 100$;
if $m_1 - m_2 = 1$ then $I_2/I_1 = 10^{0.4} = 2.512$ (using a calculator).
Both of these are what we expect them to be from the earlier discussion.

This magnitude is called the *apparent magnitude* because it describes stars as they appear in the sky. The brightness depends on the distance and we can standardise it, for stars of known distance, by defining an *absolute magnitude* which is the magnitude a star would have if seen at a distance of 10 pc. The choice of 10 pc is arbitrary.

Colour and Magnitude.
All this is based on looking at the stars with the human eye, so it depends on the response of the eye to different wavelengths. With a particular combination of filters and a photoelectric detector it is possible to define a 'visual' magnitude that responds very much as the eye does—and in a perfectly reproducible way. In this system, apparent magnitude is written m_V or V and absolute magnitude M_V. The signal detected depends on the amount of the starlight in the region defined by the filters used.

In the same way, we can define other filter-detector combinations that let us measure magnitudes in other bands, in particular blue ($m_B = B$, M_B) and ultraviolet ($m_U = U$, M_U). (*These are shown in Figure 19.8*). The three-colour photometry UBV has been widely used. The word *photometry* simply means the measurement of brightness. The system has been extended with other similar bands into the red and infrared, but these are in practice less often used.

These are 'broad' bands, each being about 100 nm wide. There are also systems of narrow band photometry, with bands typically 10 nm wide at carefully-selected wavelengths.

The difference between two of these magnitudes, such as $B - V$, gives a measure of the colour of a star. A yellow star such as the Sun will have more radiation in the V band relative to B than will a hot, blue star. A *colour index* such as $B - V$ is useful because it is independent of the distance to the star. Two stars of the same surface temperature have the same colour, and they have the same $B - V$ even if they are at very different distances and so very different values of V or of B separately.

The zero point of the V magnitude is defined by requiring that sixth magnitude stars correspond to the old visual estimates. For the other bands, we define it so that $B - V = 0$ for a white star such as Sirius, with $T_{eff} \sim 10^4$ K, and similarly for U, etc.

To think which side of this temperature has negative $B - V$ and which positive, we have to be careful because of the counter-intuitive way magnitudes behave. For the *same* V, a hotter star will have more radiation in B so B is a smaller number and $B - V$ is negative. A cooler star will have less radiation in B so B is a bigger number and $B - V$ is positive. For example, the Sun has $B - V = +0.62$.

[Here are some mathematical notes that were not given in the lecture but should be useful in understanding the subject.

We have seen earlier that the intensity is proportional to one over the distance squared (see also the discussion of Luminosity in the next lecture). That is, for distance d we have $I_d = k/d^2$ where k is some constant. Thinking of distance d and distance 10 pc, k is the same for both so

$$\frac{I_d}{I_{10}} = \frac{k/d^2}{k/10^2} = \frac{10^2}{d^2} = \frac{100}{d^2}.$$

If we call the magnitude at d, m, and the absolute magnitude at 10, M, we can use the relation

$$\frac{I_2}{I_1} = 10^{0.4(m_1 - m_2)}$$

to obtain

$$\frac{I_d}{I_{10}} = 10^{0.4(M-m)}$$

so we have the relation

$$\frac{100}{d^2} = 10^{0.4(M-m)}.$$

{For those of you who have met logarithms, taking logarithms to base 10 this can be written as

$$2 - 2\log d = 0.4(M - m)$$

or

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

which is given in the textbook. We don't use logarithms in this course.}

Applying the above relation to the V and B magnitudes, we have

$$\frac{100}{d^2} = 10^{0.4(M_V - V)},$$

$$\frac{100}{d^2} = 10^{0.4(M_B - B)}$$

so

$$10^{0.4(M_V - V)} = 10^{0.4(M_B - B)}$$

$$0.4(M_V - V) = 0.4(M_B - B)$$

$$M_V - V = M_B - B$$

$$B - V = M_B - M_V.$$

So for any distance d the colour index $B-V$ equals the difference in absolute magnitudes, $M_B - M_V$. The colour index is therefore independent of distance.]

Lecture of 5 November, 2002.

Luminosity.

The luminosity L of a star is the total energy that it radiates in 1 second, at all wavelengths. This comes out equally in all directions. The energy from unit area of the surface is σT_{eff}^4 (this is the definition of T_{eff}) and if the radius is r the surface area is $4\pi r^2$ so $L = 4\pi r^2 \sigma T_{eff}^4$.

Luminosity is usually expressed in terms of the solar luminosity $L_\odot$. Values range from about $10^6 L_\odot$ down to $10^{-4} L_\odot$ or even fainter.

Luminosity is an absolute property of the star. What we can measure is the intensity I of radiation crossing unit area of a detector at a distance d from the star. The number of unit area elements at this distance is $4\pi d^2$ so

$$I = \frac{L}{4\pi d^2} = \frac{r^2}{d^2}\sigma T_{eff}^4.$$

The wavelength dependence is similar to the shape of a black body curve but scaled down because of distance.

Spectroscopy.

This is the study of spectral lines, produces by atoms or molecules. It is by far the largest area of present-day astronomical observation.

Figure 6.23 shows a series of absorption lines imposed on a continuous emission spectrum such as the emission from a star. In the region of a spectral line, there is a reduction in the intensity of the light and it looks dark across the spectrum. A line can have different strength of absorption and it has a distinct width. By analysing the form of the line profile we can get information about physical conditions in the source.

We can also have emission lines, where there is an excess of intensity over the continuum. *Figure 27.3* shows an example of this; in fact this is the spectrum of a quasar. It is possible to have emission lines even if there is no continuous emission.

Atoms.

An atom is somewhat like a miniature planetary system. Almost all the mass is in the central nucleus, about 10^{-14} m in diameter, and circling round it in orbit are several electrons. The orbits are about 10^{-10} m or 0.1 nm in radius. The atom is held together by electrical forces. In electricity there are two types of charge, which arbitrarily are called positive and negative. In the atom, the electrons are identical and each has the same negative charge. The nucleus has a positive charge equal in magnitude to the sum of the charges on the electrons, so overall the atom is electrically neutral. It is possible for an atom to lose one or more of its electrons and become a *positive ion*; in some cases an extra electron can be attached to an atom to form a negative ion.

[*See Box 5.5*]. The Periodic Table is structured in a way that relates to the chemical properties of the atoms; we can use it as a convenient list of elements each with its name, chemical symbol and atomic number (the number of electrons and the charge on the nucleus in units of the electron charge).

The various elements have different abundances in astronomical objects; to a good approximation these cosmic abundances are the same everywhere (shown *approximately* in *Figure 7.11* as relative numbers of atoms, relative to 10^{12} for hydrogen (H)). The most important elements for us are hydrogen (H), helium (He), carbon (C), nitrogen (N) and oxygen (O). Details in the relative abundances arise from the processes inside stars by which the elements are formed and destroyed.

Lecture of 12 November, 2002.

The Hydrogen Atom.

Hydrogen is the most abundant atom by far, so it is very important in astronomy. It is also the simplest possible atom, with just one orbiting electron and a nucleus with charge equal to the electron charge. The hydrogen nucleus is called a *proton*.

The quantum theory of the hydrogen atom was given by Bohr in 1912. The essence is that the electron can go round only in certain particular orbits. each with its own energy. The atom absorbs or emits radiation when the electron jumps between two orbits. When it jumps to an orbit of greater energy, radiation is absorbed; when it jumps to an orbit of lower energy, radiation is emitted. The radiation is in the form of a photon, of energy $h\nu$ equal to the energy difference between the two orbits.

The orbits are labelled by a *quantum number* n, $n = 1$ being the lowest energy orbit. Radiation comes when the electron jumps between two orbits n_1 and n_2. It is absorption or emission according to which way the electron jumps. (*See Figures 5.20, 5.21*).

[*Not in the lecture.*
The radius of an orbit is proportional to n^2; the energy of the atom when the electron is in orbit n is proportional to $-1/n^2$. When n is very large the atom has zero energy which corresponds to the orbit being very large. The wavelength of the transition between orbits with $n = n_1$ and $n = n_2$ where n_1 is the inner orbit of the two is in this theory given by

$$\frac{1}{\lambda} = R\left(\frac{1}{n_1^2} - \frac{1}{n_2^2}\right)$$

where R is a number called *Rydberg's constant.* For instance, if $n_1 = 2$ and $n_2 = 3$ then

$$\frac{1}{\lambda} = R\left(\frac{1}{2^2} - \frac{1}{3^2}\right) = R\left(\frac{1}{4} - \frac{1}{9}\right) \ ]$$

With $n_1 = 2$ and higher values of $n_2 = 3, 4, 5, ...$, we get different energies and wavelengths and this forms a series of spectral lines in the visible, called the Balmer series:

$n_1 = 2$	$n_2 = 3$	656 nm	Red	Hα	or Balmer α
	$n_2 = 4$	486 nm	Green	Hβ	
	$n_2 = 5$	434 nm	Blue	Hγ	
	$n_2 = 6$	410 nm	Violet	Hδ	
	.....				

With other values of n_1 we get other series of lines. $n_1 = 1$ gives the Lyman series, in the ultraviolet. This is conveniently shown in an energy level diagram (*Figure 5.23*).

If the atom absorbs energy that takes it above $n = \infty$ in the diagram, the electron is totally removed from the atom and can escape. We end (in this case) with a bare nucleus and a separate, free electron. This process is called *ionization.*

The energy levels are often referred to as the different *states* of the atom. For hydrogen, the $n = 1$ level is the *ground state* and in general this is the name given to the lowest-energy state of any atom.

Other Atoms.

Any atom has a nucleus and orbiting electrons. For hydrogen the one electron can be only in certain orbits, with particular values for the energy of the atom. For an atom with several electrons, each electron can be only in certain possible orbits and the total energy of the atom can take only certain particular values. There is an energy level diagram and a set of spectral lines, for each atom. The lines give a signature for that atom; if we see lines belonging to an atom in the spectrum of a star we know that the atom is present there (*See Figure 5.15*). We can find the relative abundances of different atoms (with some interpretation—each line has its own intrinsic strength, and the relative strengths of lines can depend on the temperature).

If we remove an electron from an atom, to produce a positive ion, there is a totally different set of spectral lines, and so on with the successive ions as electrons are removed, until there is a bare nucleus.

Hydrogen has one electron and so just one spectrum.

Helium has two electrons.

He I	is the spectrum of	He	(2 electrons)
He II		He^+	(1 electron)

Carbon has six electrons.

C I	is the spectrum of	C	(6 electrons)
C II		C^+	(5 electrons)
C III		C^{2+}	(4 electrons)
C IV		C^{3+}	(3 electrons)

and so on. When carbon is vaporised in the laboratory, its spectrum is seen and is called C I. Increasing the temperature, these lines fade out as the atoms become ionised and a second set of lines appear, called C II, and so on as it is heated still more.

In a cool stellar atmosphere the carbon atoms are mostly neutral. In one slightly hotter the mixture of C^+ relative to C increases so the C I lines are weaker and C II lines stronger. For a hotter star still, the amount of C^{2+} increases so the C II lines become weaker and C III lines appear. This happens for all atoms. In the case of hydrogen, in hot stellar atmospheres the hydrogen is mostly ionised so the hydrogen lines are weaker.

The observed strengths of optical lines such as the Balmer lines in H also depends on temperature in a different way. The Balmer lines are absorptions from $n = 2$. At low temperatures there is not much energy available and most of the H is in $n = 1$. As temperature increases, the $n = 2$ population grows and the Balmer absorption lines are stronger. For higher temperatures the H becomes mostly ionised, H^+, there are fewer atoms and the Balmer lines are weaker again. So there is a certain temperature for which the Balmer lines are at their strongest, which is in fact about 10,000 K. In general, for each atom and ion there is a temperature for which the optical lines at at their strongest (*See Figure 19.12*).

This gives a measure of the temperature T_{eff} by comparing the ratios of the observed strengths of suitable spectral lines.

Lecture of 19 November, 2002.

Spectral Classification.

The system used to classify stars from their spectra was introduced at Harvard early in the 20th century. It started by labelling A the stars with the simplest-looking spectra, B for the next simplest, and so on. Later it was realised that spectral appearance follows a temperature sequence and (with some changes in the classes included) the modern sequence is OBAFGKM from hottest to coolest. Sometimes classes RNS are added; these are stars with unusual abundances that parallel K and M in temperature. Classes are subdivided into ten, as G0 to G9 (O stars have 5 to 9 only). The Sun is of class G2.

The spectral class corresponds quite closely to the effective temperature and also to the colour index $B - V$. These relations are not exact as the different quantities can be affected differently by other factors such as the radius and luminosity of the star.

Fraunhofer Lines.

The dark lines in the solar spectrum were first noticed by Wollaston, who repeated Newton's prism experiment (*Figure 5.4*) using a hole in the form of a narrow slit lined up with the dispersion direction. This gives a better spectral resolution and some of the strong absorption lines are seen. These lines were studied in detail by Fraunhofer, early in the 19th century, who catalogued them, attaching letters to the strongest features. Later in the century it became possible to identify the lines with laboratory spectra of various atoms.

Nowadays only a few of Fraunhofer's designations are in common use:
D: Two lines of Na I (sodium), close together in the yellow.
H and K: Two lines of Ca II (Ca^+, ionised calcium), rather less close together in the violet. The Ca II H line is blended with hydrogen $H\epsilon$ in the spectrum of a star.

The Doppler Effect.
It is found that if the source of a wave and an observer are moving relative to each other then the wavelength shifts. This is most familiar with sound waves; as a moving vehicle speeds past the pitch of the sound drops. Pitch is measured by the frequency of the wave; for motion towards us the frequency is higher than at rest, for motion away it is lower.

This happens for light too. Motion towards the observer means a shift to shorter wavelength (a blue shift), motion away a shift to longer wavelength (a red shift). We can understand this qualitatively—moving away stretches the wave so the wavelength is longer. (*See Figure 5.24*).

The effect happens at all wavelengths but we can see it and measure it only using the spectral lines. The Doppler shift formula is that, if the observed wavelength is λ (lambda) and the wavelength at the source (the rest wavelength) is λ_0 then

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

where v is the speed of the source relative to the observer and c is the speed of light. The same formula holds for all wavelengths. v is the speed in the line-of-sight direction, the *radial velocity.* If λ is larger than λ_0, v is positive. So v is positive for motion away from the observer, negative for motion towards the observer.

The random motions of nearby stars relative to each other and to us are of order 20 km sec^{-1}, with both positive and negative values of v. The galaxies within about 1 Mpc distance are members of our Local Group, which is a small cluster held together by gravity. There are more than 40 known members, most of them very small but including also some large galaxies such as the Milky Way galaxy itself or the Andromeda galaxy M31. The Local Group galaxies are moving around at random and have both positive and negative radial velocities up to a few hundred km sec^{-1}. (*See Figure 26.18, which is a projection of the three-dimensional distribution onto a plane*).

For galaxies outside the Local Group, there is a systematic motion away from us that increases with distance. This is the expanding Universe and it is expressed in Hubble's Law $v = H_0 d$ where H_0 is a constant. We can define a quantity called the *redshift*,

$$z = \frac{\lambda - \lambda_0}{\lambda_0}.$$

Galaxies are now known to $z = 1.0$ or 2.0; the most distant known quasars have z greater than 6.

From this, $\lambda/\lambda_0 = z + 1$. As an example, for $z = 4.0$ the Lyα line is shifted from its rest wavelength of 121.6 nm, far in the UV, to $121.6 \times 5 = 608$ nm, in the red part of the visible.

The formula

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

applies if v is much smaller than c. If v becomes a significant fraction of c then a different form of the equation applies, that takes Relativity into account. For such values, z is no longer given by v/c but instead by a relativistic formula,

$$z = \sqrt{\frac{c+v}{c-v}} - 1, \quad \text{or} \quad \frac{v}{c} = \frac{(z+1)^2 - 1}{(z+1)^2 + 1}.$$

Thus, if $z = 4$ then $v/c = 24/26 = 0.923$. v can never be greater than c.
[Note that the textbook (page 620) quotes a formula for v/c that is not correct. The example given uses the correct relation.]

Proper Motion.
The Doppler shift gives the radial or line-of-sight component of a star's motion. There is also a transverse component or *tangential velocity*. Radial velocity depends on the shift of spectral lines so it can be measured for objects at any distance, even the most distant objects in the Universe. Tangential velocity depends on our ability to measure the change in direction of an object in the sky. (*See Box 19.1*).

The *proper motion* is the angular shift in one year. Barnard's star has the largest known proper motion, 10.4 arcsec y^{-1}. Proper motions can be measured only for relatively nearby stars. The *Hipparcos* satellite helped because of its accuracy but it is also important to use a time baseline of many years. For galaxies and quasars we can measure only the radial velocity.

Binary Stars.
More than half the stars we see in the sky are binary or multiple. A single isolated star like the Sun is not unusual but they are in the minority. A binary star has two components that are close enough to influence each other. Each component is an ordinary star with its own spectral class although if they are very close they can modify the physical properties. We can observe binaries in three different ways.

1. Visual binaries.
The stars are far enough apart that we can see them separately, and see the motion of one round the other, or rather of each round the centre of mass of the system (*See Figures 19.19 to 19.21*). This distinguishes the binary from an optical double, a chance alignment of stars at very different distances.

If there is proper motion, the centre of mass travels in a straight line and each star shows a wobble about this line. The star Sirius has a very faint white dwarf companion; its presence was deduced from the wobble in the motion of the primary star long before it was seen directly.

From the detailed form of the orbit, information can be obtained about properties of the two component stars.

2. Spectroscopic binaries.
The two component stars are so close together that their orbital velocities are large enough to give measurable Doppler shifts. (*See Figure 19.24*). This means that they are too close to be seen separately so a spectroscopic binary is not also a visual binary.

We can plot the radial velocity curves as a function of time. Very often only one set of spectral lines can be seen, moving back and forth periodically, in what is called a single-line binary: this happens if the secondary component is more than a few magnitudes fainter when its light is swamped by the primary.

3. Eclipsing variables.
If one star passes in front of the other it blocks out some of the light and the total intensity falls. This can happen more readily if the stars are close together and in practice eclipsing binaries are also spectroscopic. Depending on the geometry of the system, the light curve can take different forms and by analysing it and the velocity curve we can find out a lot about the stars and their orbit. (*See Figures 19.26, 21.17*).

It is from observations of binary stars that we get measurements of stellar masses: there is no way to measure the mass for a single isolated star.

Lecture of 3 December, 2002.

Intrinsic Variable Stars.
Eclipsing binaries are one class of variable star, where the intensity received varies as a function of time. There are also some types of single stars where there is an actual variation of the light produced. They are classified in a variety of types.

Some have totally irregular variations, caused by some process affecting the surface of the star. An example is the T Tauri stars (named after a typical example—the use of one or two capital letters with the constellation name denotes a variable). These are newly formed stars still on the way to the main sequence, that have not yet settled down to shine steadily.

Some stars are physically pulsating, changing in radius in a periodic way. Mira (*Figure 21.12*) is a long-period variable with a fairly regular period of almost one year.

The Cepheid variables (*Figure 21.14*), named after the typical example δ (delta) Cephei, have very regular and repeatable behaviour. The brightness and radial velocity vary with the same period. This comes from a pulsation in which the radius changes by 15% (for δ Cep; in other cases it can be larger). The Cepheids have important properties that make them useful in distance determination.

Stars with Emission lines.
The spectral lines normally seen in stellar spectra are absorption lines. These come from the visual surface of the star, called the *photosphere.* Emission lines can come from a hotter outer region. For the Sun, the corona is at high temperature (about 2×10^6 K) and very low luminosity. The amount of material is small and the amount of radiation is much less so that the corona can't be seen at the same time as the photosphere. The corona has an emission line spectrum; the strongest line, at 530 nm, comes from Fe XIV, that is iron that has lost 13 of its 26 electrons—characteristic of the high temperature.

Some other stars have extended outer shells which give emission lines that are strong in relation to the photosphere and its absorption lines so they can be seen in a spectrum of the star. In some cases the material in the shell may be falling in or expanding out; this can be seen as a Doppler shift relative to the photospheric lines.

Nebulae.
A nebula is an extended object (that is, not point-like like a star) that is fixed among the stars and so is outside the Solar System. The main catalogues are that of Messier (1781) and the New General Catalogue (1888) and the M or NGC numbers are used to refer to these objects. These lists contain objects of different nature, both gas and dust clouds within our Galaxy, and other galaxies some at very great distances. These were distinguished only when large telescopes could show their shapes, in particular the spiral or regular elliptical shapes of galaxies compared with the irregular interstellar clouds.

A galaxy has a spectrum much like an average star, with continuous emission from all the stars combined and absorption lines that are among the strongest of the stellar lines (such as Ca^+ H and K); the weaker lines are smeared out by differences of spectral type and of Doppler shift. By contrast the gaseous nebulae have emission lines.

Emission Nebulae.
The gaseous nebulae typically have temperatures around 10^4 K and densities 10^9 to 10^{10} atoms m^{-3}. There are three types, according to their situation.

1. H II Regions.
The ultraviolet radiation from O and B stars has sufficient energy to ionise hydrogen atoms and to produce a region around the star where H is predominantly in the ionised form H^+ (typically 95% or more). In an extension of the notation used for spectra, this is called an H II region. Outside it, the hydrogen is neutral (H I region).

The shape is generally very irregular because of density fluctuations. These OB stars are all young and are close to the interstellar clouds from which they formed (unlike an old star like the Sun). A good example is the Orion Nebula (*Figure 20.1*).

The emission lines are generally dominated by the hydrogen Balmer lines. Hα is the strongest, giving most of these nebulae their characteristic red colour.

2. Planetary Nebulae.
After a star has been a red giant it throws off a lot of its mass. From being a shell around the star this material develops into something with properties similar to H II regions but rather more regular in shape. (*See Figure 22.6 and the cover of the textbook.*)

3. Supernova Remnants.
When a massive star explodes, it throws off a lot of material which collides with the ambient interstellar material and gives a shell-like structure. The spectrum is similar to those from H II regions and planetary nebulae but in this case the energy radiated originated in the explosion rather than coming as radiation from the hot central star or stars.

The typical lifetime of a planetary nebula or supernova remnant is around 10^4 years, after which the material merges into the general interstellar medium.

Examples of nearby supernovae include 1987A in the Large Magellanic Cloud (a satellite galaxy of the Milky Way, in the southern sky) (*Figure 22.16*), and the Crab nebula which is seen where a supernova was recorded in AD 1054; the Crab itself was first recorded by Lord Rosse.

Remnants of older supernovae include the Veil Nebula in Cygnus (*Figure 22.22*), known to radio astronomers as the Cygnus Loop.

Reflection Nebulae.
There can be bright blue nebulosity as well as red (*See Figures 20.4, 20.18*) sometimes in the same region (*See page 474*). The blue light is a reflection of starlight; it can happen near hot stars but outside the H II region, or near cooler stars that don't have enough UV radiation to ionise hydrogen but are still reasonably hot, such as A stars.

The reflection is by the grains of interstellar dust, which reflect much more efficiently than the gas does. These are small solids, less than 1 μm in size. Blue light is reflected more than red, so what we see looks blue. The amount of reflection varies with wavelength approximately as $1/\lambda$. (For the Earth's atmosphere, where the scattering that produces blue in the sky is mainly by molecules, the wavelength dependence is proportional to $1/\lambda^4$ so again blue is scattered more than red).

The light from a star that has suffered this scattering, seen directly, looks redder than it would have done. (*See Figure 20.5).* It is also fainter overall, so the process is called both *interstellar extinction* and *interstellar reddening.*

Extinction can blot out completely the visible and UV radiation from stars beyond 2 or 3 kpc in the plane of the Galaxy.

Lecture of 10 December, 2002.

Cold Interstellar Gas.
We can see dark nebulae in silhouette against bright material (stars, and emission or reflection nebulae). (*See Figure 20.2, 20.3*). This is the result of obscuration by the dust, which is mixed with cold interstellar gas. Overall, rather a small fraction of the interstellar material is in the bright emission nebulae, which are very localised, or in these dark clouds seen by chance in silhouette.

The general interstellar gas has a mean density of about 10^6 atoms m^{-3} and temperature around 100 K. It can be seen through absorption lines in starlight, which are much narrower than stellar absorption lines (a result of the lower density). This is naturally only for stars that can be seen through the obscuration; the dust that is mixed with the gas totally obscures the more distant stars.

The dust itself can be observed as a thermal continuum emission in the infrared (*See Figure 25.7*). The dust temperature is typically 10–20 K but close to stars it can be warmed to as much as 90 K.

Radio waves can penetrate through the dust and there is a spectral line at 21 cm from interstellar hydrogen that can be used to map the whole Galaxy, and other galaxies too. This comes about because in the H atom the proton and the electron have magnetic properties that can be described in terms of a spin around an internal axis. The two spins can be parallel or opposite (*Figure 25.10*), with no other possibility. These two situations correspond to very slightly different energies. There is thus a splitting of the $n = 1$ state in the H atom into two energy levels. The transition between these gives radiation of wavelength 21.1 cm.

The 21 cm line is widely seen in emission from interstellar gas (which is predominantly hydrogen). Mapping the plane of the Milky Way (*Figure 25.13*) shows a structure that by analogy with other galaxies can be interpreted as a spiral pattern.

Structure of the Galaxy.
(Note the convention that we use capital G for our Galaxy, lower case g for others.) (*See Figure 25.8, 25.15*). There is a flat disk, a few hundred pc thick and 25 kpc or more in radius. This contains most of the interstellar gas and dust, and young stars many of which are in clusters of 100–1000 members. These are called open clusters or galactic clusters (the latter name therefore shouldn't be used for clusters of galaxies).

The Sun is in the central plane, about 8 kpc from the Galactic Centre. It is part of a general distribution in the plane of older stars both inside and between the spiral arms.

There is a central bulge region, spherical in shape and a few kpc in radius containing mostly older stars and not so much interstellar matter. There is a halo, essentially an outer spherical distribution, containing some stars and interstellar matter, but not much of either, and the globular clusters which are older than open clusters and contain about 10^4 stars each.

The Galaxy is rotating. The Sun and neighbouring stars are together going round at about 250 km s^{-1}. The spiral arms trail.

Other Galaxies.
In the late 19th and early 20th centuries there was a great argument about how far away the nebulae are. It was settled when the 100-inch telescope on Mt Wilson came into use and was able to resolve individual stars in other galaxies, proving that they are distant. Part of the argument had been that there appear to be fewer galaxies near the plane of the Milky Way. This is caused by obscuration by interstellar dust in the plane of our Galaxy; the presence of the dust was not conclusively established until about 1930.

Classification.
We classify galaxies according to their appearance in the sky. The basis of the classification used was given by Hubble and we look at it here in outline.

1. Spiral galaxies. (*Figure 26.4, 26.5*)
These are essentially like our Galaxy, although there are variations in the detailed shapes. Visible masses range from about 10^{11} $M_\odot$ downwards, about 10% of which is interstellar matter. (Possible unseen dark matter could increase the galaxy mass by as much as a factor of 10).

2. Elliptical galaxies. (*Figure 26.6. Note that for M105 the wrong picture is given—an E0 galaxy should be circular*).
The shape in the sky is of an ellipse, a flattened circle. Masses range from about 10^{12} $M_\odot$ down to dwarf ellipticals of 10^6 $M_\odot$ or less. There is much less interstellar matter

than in spirals, generally less than 0.1% of the mass. Corresponding to this there are few young stars—the population is like the bulge of our Galaxy or a globular cluster.

3. Irregular galaxies.
These have masses from about $10^7\ M_\odot$ down and about 25% of the mass is interstellar. There are many young stars. Examples include our satellites, the two Magellanic Clouds (*Figure 26.10, 26.11*).

Clusters of Galaxies.
Galaxies tend to be in clusters (*Figure 26.7, 26.17, 26.20*). The richest clusters have more than 1000 known members. The separations of the galaxies in a cluster are much smaller than the distance to the cluster so they are effectively at the same distance.

Active Galaxies.
These are normal galaxies with an active nucleus at the centre. This is understood to be caused by the presence of a massive black hole there, of mass perhaps $10^{10}\ M_\odot$. Matter flowing in to the black hole releases a great amount of energy which is seen as non-thermal radiation (that is, radiation not produced by normal processes involving the temperature such as for stars). The radiation is particularly seen in the radio and in X-rays. Some of the radio sources are very extended.

Seyfert galaxies (*Figure 27.9*) are a category where the active nucleus is in a galaxy that can be seen, and is a spiral galaxy. Quasars (*Figure 27.7*) were originally seen as strong radio sources associated with point-like optical objects, like stars (hence the name). Similar objects that are radio-quiet were later found. In recent years it has become possible to see the surrounding galaxies for a number of them, surprisingly many of which turn out to be ellipticals suggesting that the Seyferts and quasars are not simply similar objects with different relative strengths and at different distances.

The optical spectra contain emission lines, which for quasars are at redshifts that in a few cases are as large as $z = 6$ or more.