PHAS 1102
Physics of the Universe

2 – Atomic structure

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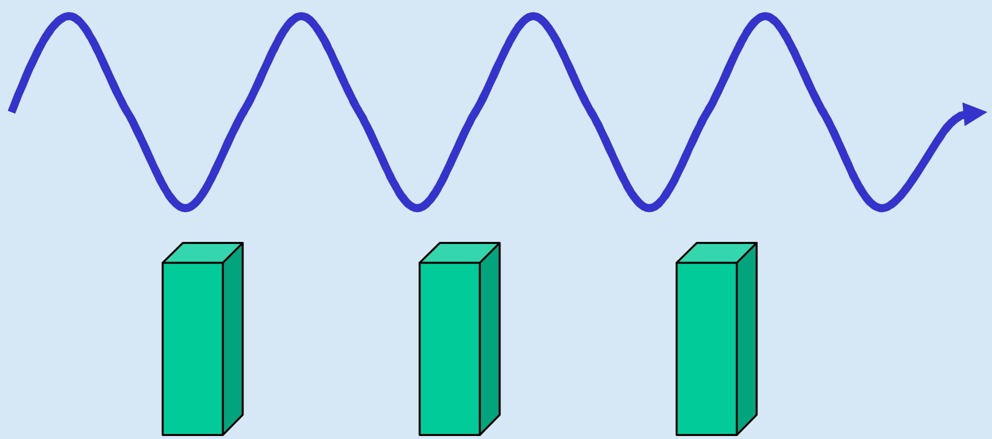
Atomic structure *(ZG Ch. 8; FK Ch. 5)*

When an electric charge is accelerated, electromagnetic energy is produced. This energy can be thought of as propagating as a wave – or, equally, as a particle:

wave-particle duality

The waves are usually referred to as **light waves** or **radiation**.

The particles are known as **photons** or **quanta of light**.



Radiation

Wave nature of light

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

λ: wavelength
c: speed of light (3×10^8 m s^{-1})
ν: frequency

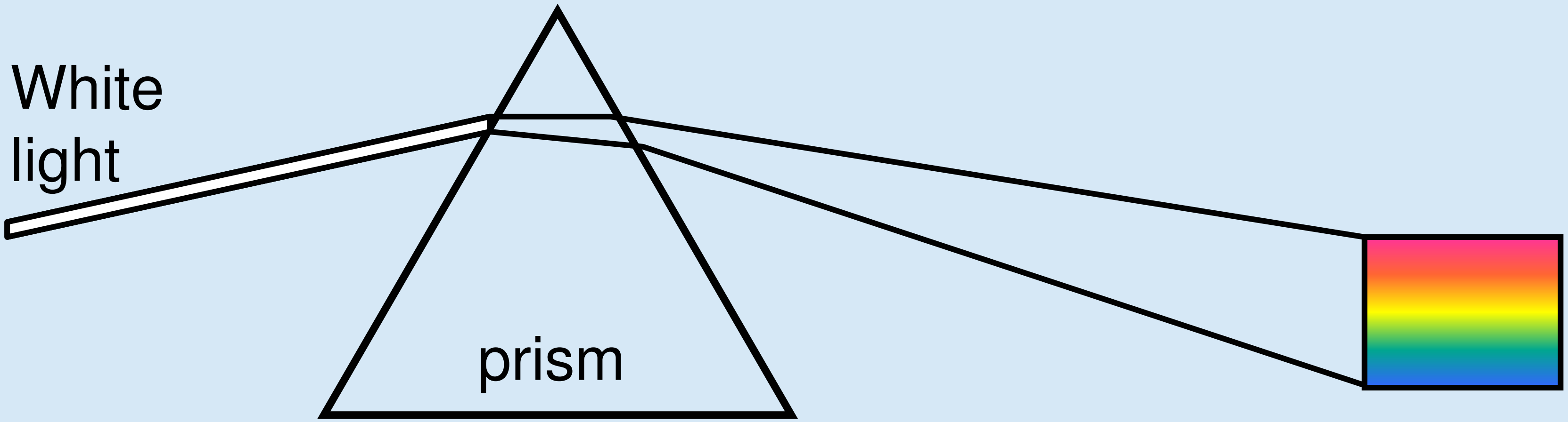
Quantum nature of light - Photons (or quanta)

$$\boxed{E = h\nu}$$

E: energy
h: Planck constant
ν: frequency

Spectroscopy

Spectra were first seen due to the effects of refraction – i.e. the bending of light as it passes through a transparent medium, which is wavelength-dependent ⇨ dispersion

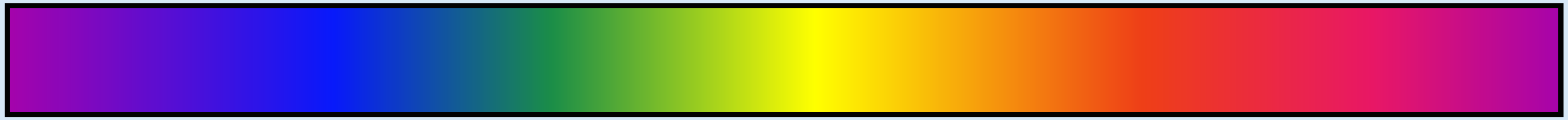


Blue light is bent the most, red light the least.

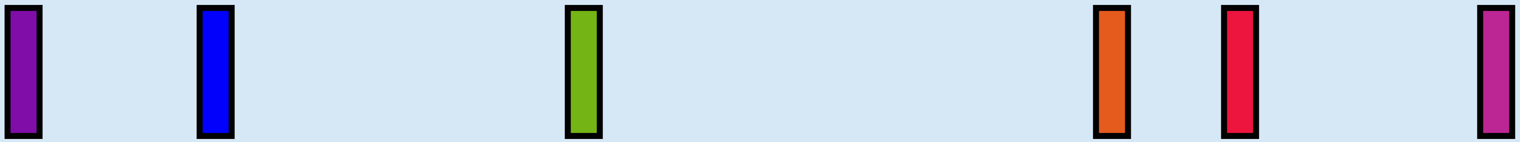
Spectroscopy is the astrophysical technique which is <u>key to our understanding</u> of astronomical objects: physical state, chemical composition, dynamics, …

Kirchhoff's (1824-1887) Rules (empirical)

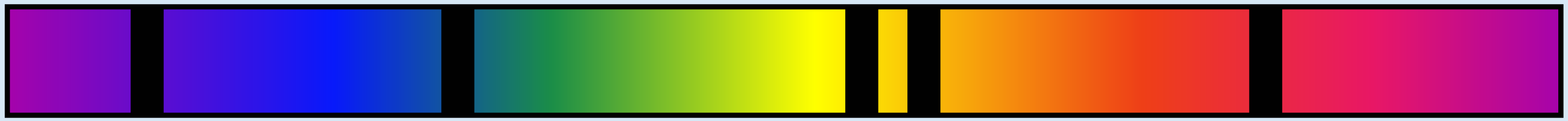
1. A hot, opaque object emits a continuous spectrum



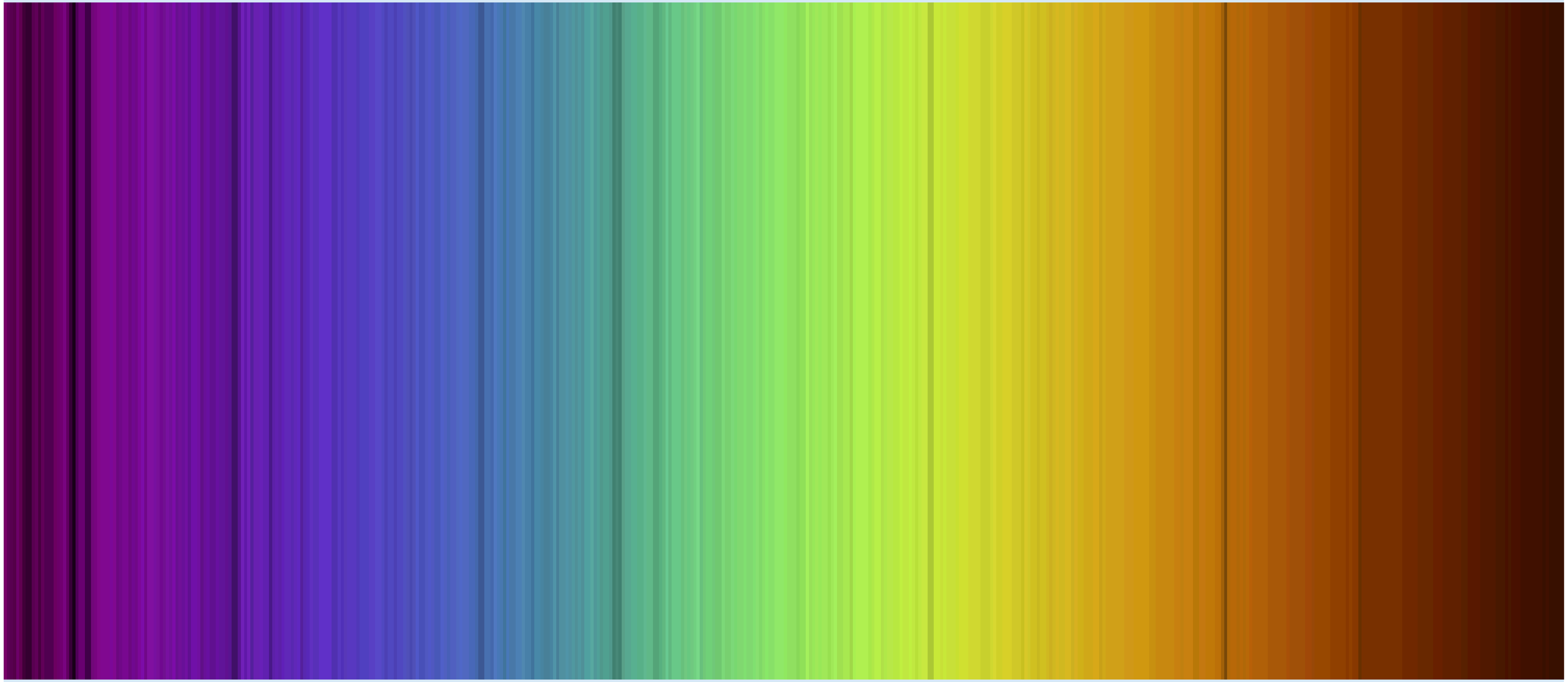
2. A hot, transparent gas produces a spectrum of emission lines. The line positions depend on the elements in the gas.



3. If light with a continuous spectrum passes through a transparent gas at a lower temperature, the cooler gas will superimpose dark absorption lines on the spectrum.



The solar spectrum



Absorption and emission lines - Atoms

Where do absorption and emission lines come from? From the absorption or production of energy when an electron in an atom changes its level (orbit).

An atom is made up of a very small nucleus, consisting of *neutrons* and *protons*, surrounded by a cloud of *electrons* (out to a radius of ~10^{-10}m) – *neutrons*, *protons* and *electrons* are elementary particles, building blocks of matter.

Proton mass = 1.6725×10^{-27} kg charge = +e

Neutron mass = 1.6748×10^{-27} kg charge = 0

Electron mass = 9.1091×10^{-31} kg charge = -e

(e = 1.6×10^{-19} Coulomb)

Bohr's atom → Quantum mechanics

In 1911 Rutherford put forward the nuclear theory of the atom with electrons orbiting the nucleus

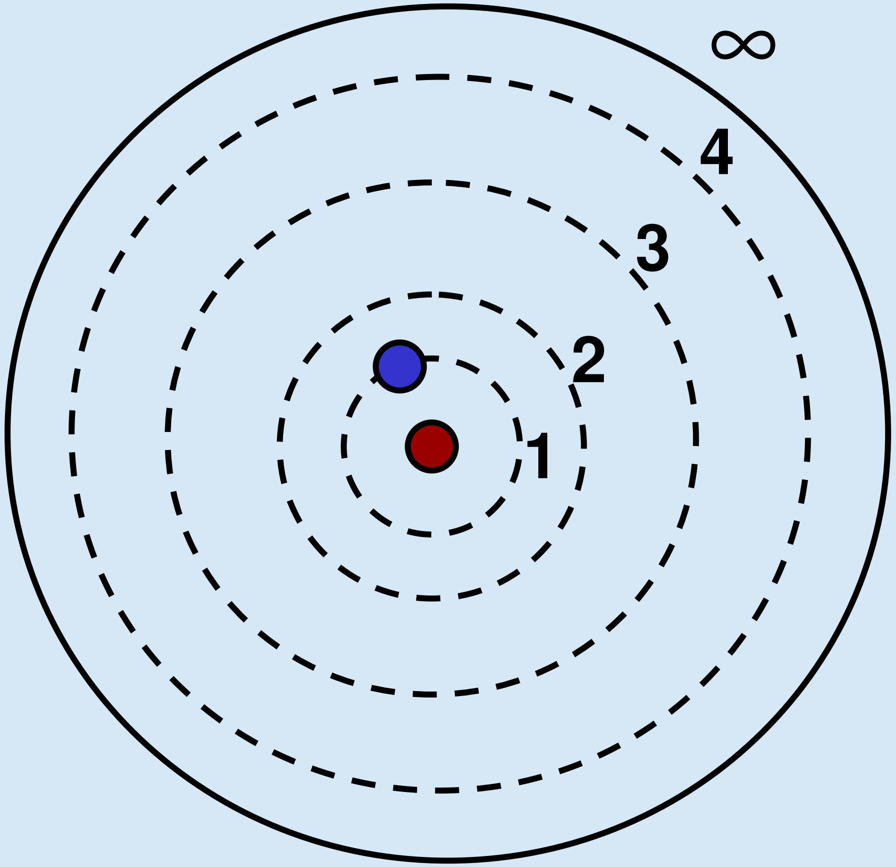
To explain the stability of this model, in 1913 Niels Bohr postulated that only a discrete number of orbits are allowed, and that in them the electron cannot radiate

Permitted orbits: those where electron's orbital momentum is an integer multiple of $\frac{h}{2\pi}$ where h is Planck's constant

→Principal quantum number n = 1, 2, 3, …

Bohr also postulated that
- a single discrete quantum of radiation is emitted or absorbed as the electron jumps from an orbit to another, and
- the energy of the radiation equals the orbits' energy difference

The hydrogen atom



Take the simple case of Bohr's hydrogen atom.

It has one proton in the nucleus and one orbiting electron.

In its stable state, the electron orbits in the <u>energy level</u> defined by n = 1 (the ground state). There are an infinite number of discrete energy levels, converging to $n = \infty$, the ionization potential.

Quantum numbers

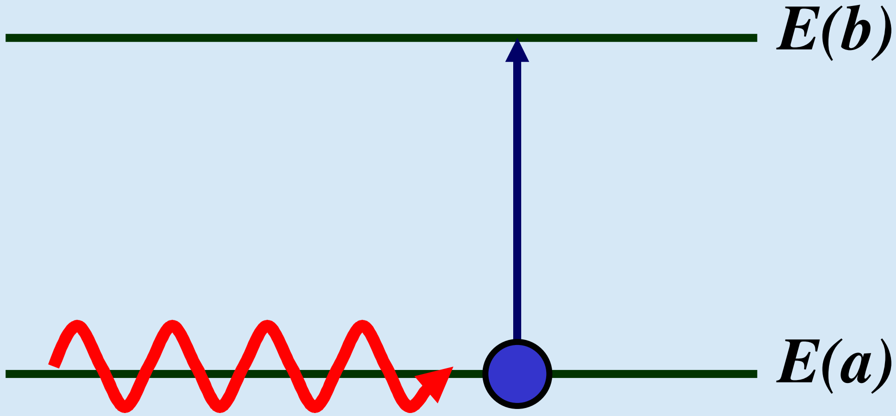
Only certain discrete energy levels are allowed for electrons in atoms. The levels are defined by quantum numbers, of which n is the 'principal' (but there are several others …).

Quantum mechanics

The electron energy levels are negative (orbits are 'bound'):

$$\boxed{E(n) \propto -\frac{1}{n^2}}$$

We can normalise to a positive energy scale by subtracting $E(1)$ from all $E(n)$.

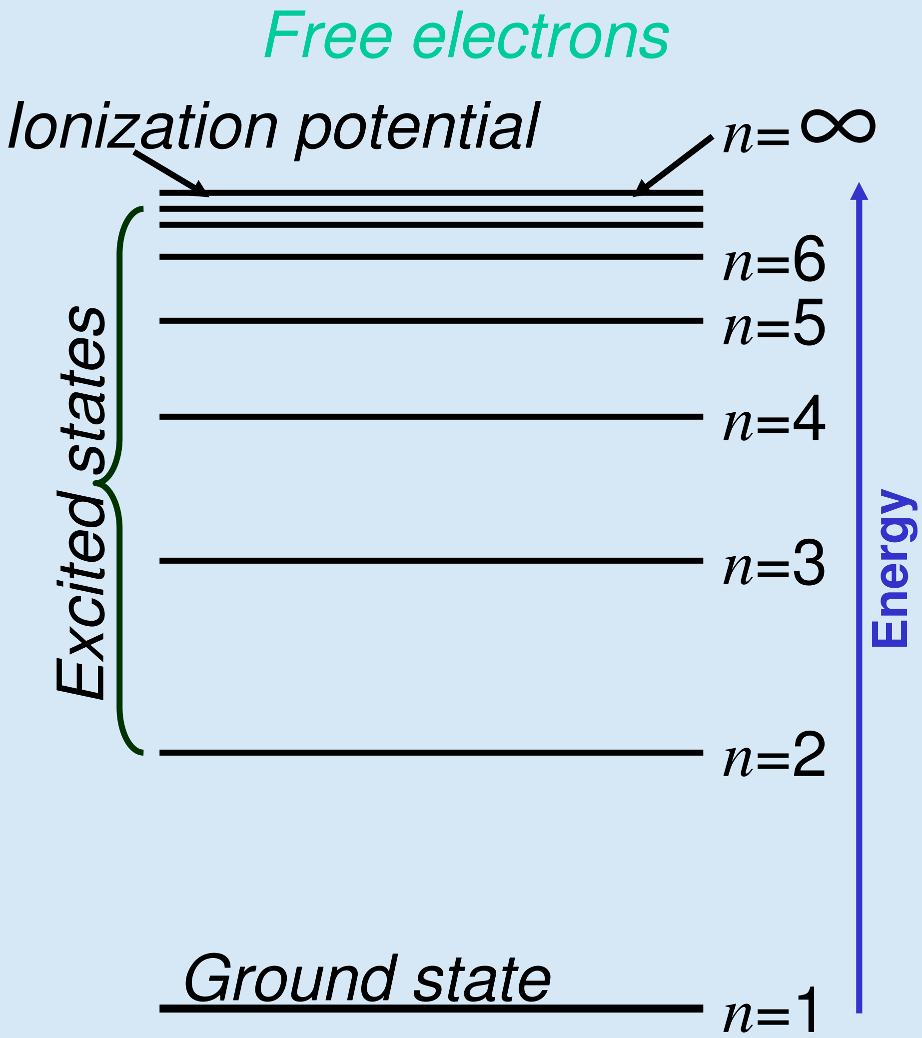


Photon energy, $h\nu = E(b) - E(a)$

$E(a)$ more negative than $E(b)$, and $E(\infty) = 0$

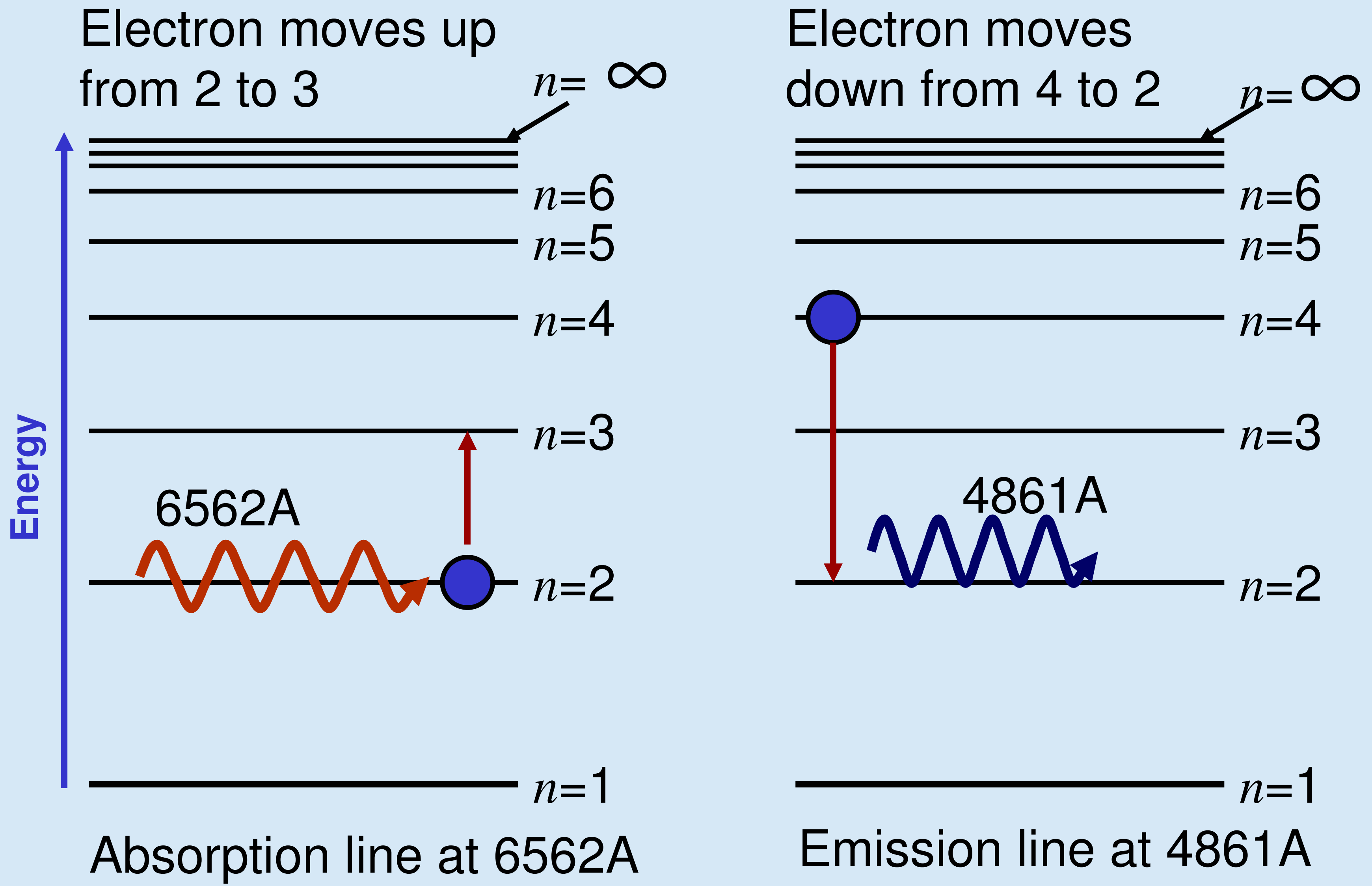
Energy levels: ionization potential

Atoms have an infinite number of energy levels, converging to a finite energy value (the ionization potential). If an electron gains more energy than the ionization potential, then it is no longer bound to the atom. Only the lowest level (the ground state) is generally stable. Excited states (when an electron is in level n = 2 or higher) have lifetimes of ~10^{-8} seconds.

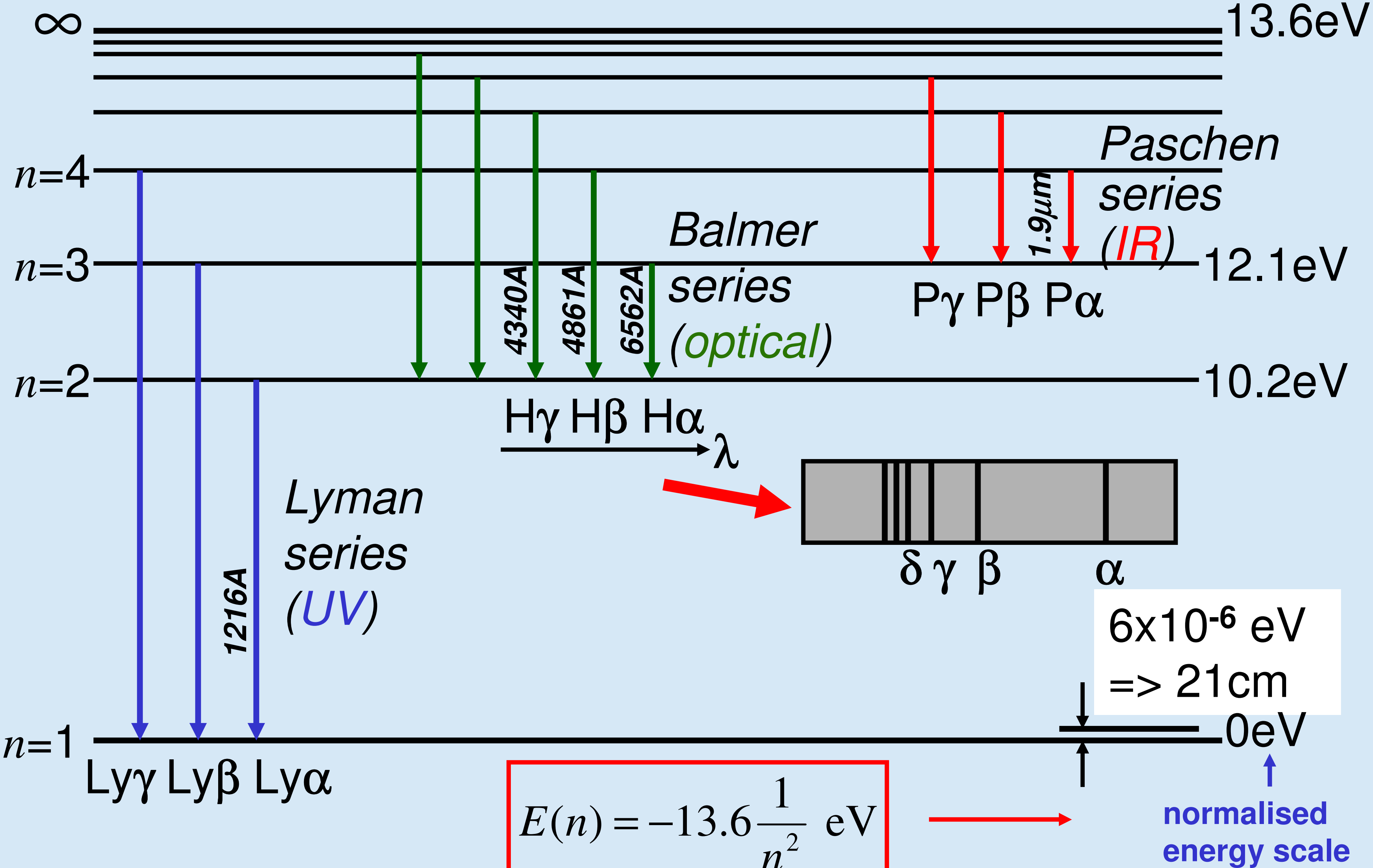


Levels structure complex, especially for atoms heavier than H: more quantum numbers exist, and many more energy levels too ...

Transitions of the hydrogen atom



Spectrum of the hydrogen atom



$$E(n) = -13.6\frac{1}{n^2}\ \text{eV}$$

Emission lines

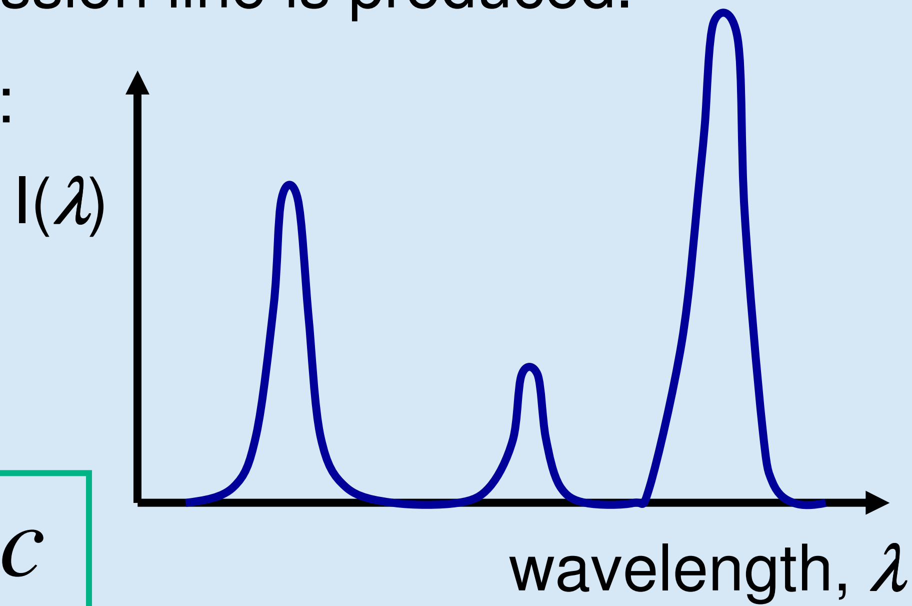
To produce emission lines, an excited state must first be populated – when the electron in an excited state falls by one or more levels, an emission line is produced.

To populate the excited levels:

(a) Collisional excitation

(b) Photo-excitation

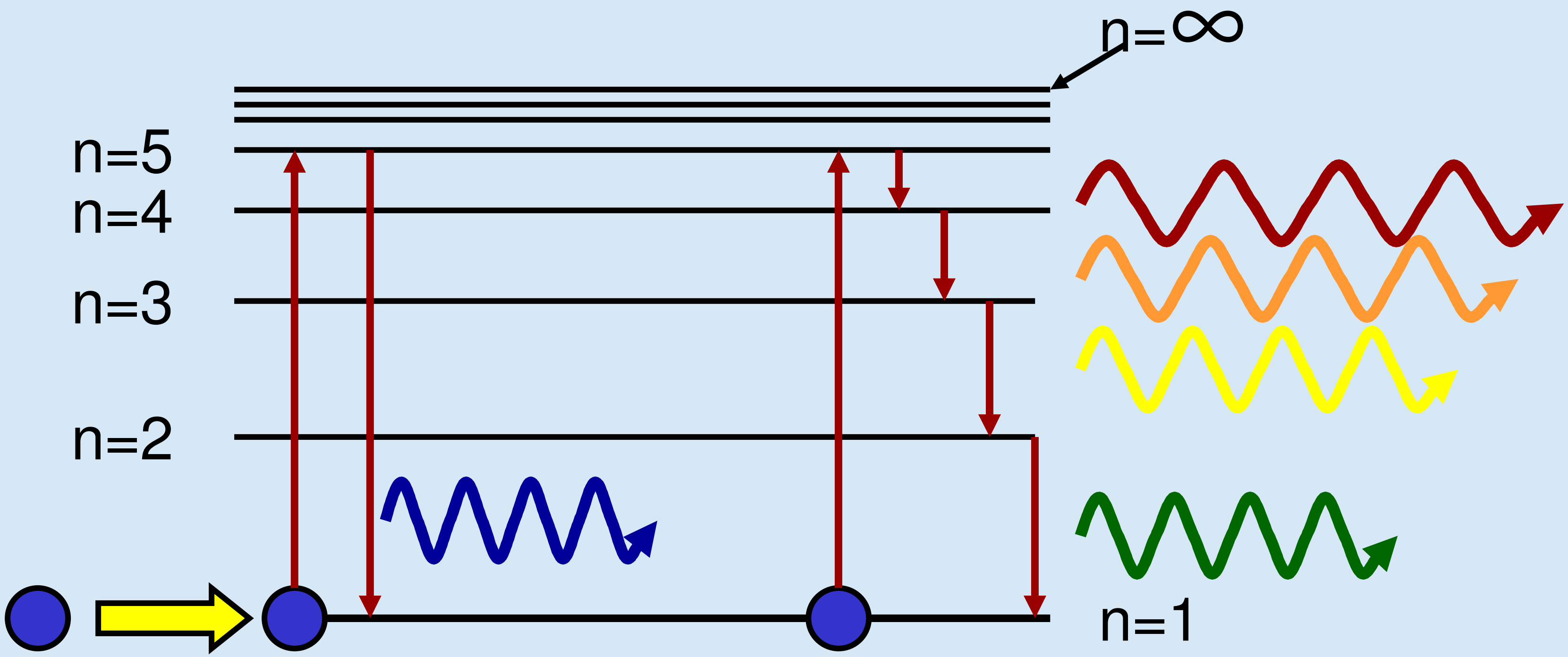
(c) Recombination



$$\boxed{\Delta E = h\nu} \Rightarrow \boxed{\lambda = \frac{hc}{\Delta E}}$$

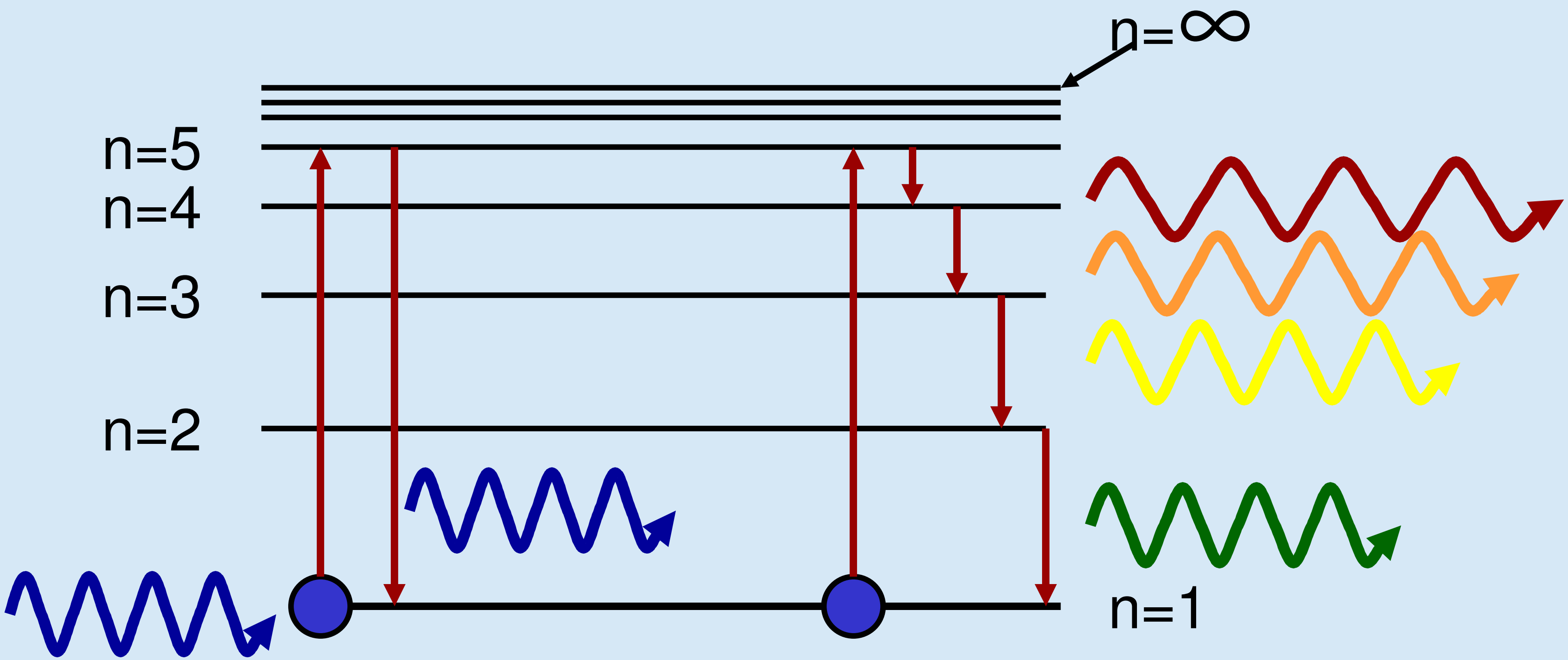
These all produce emission lines and explain Kirchoff's 2nd rule.

Collisional excitation



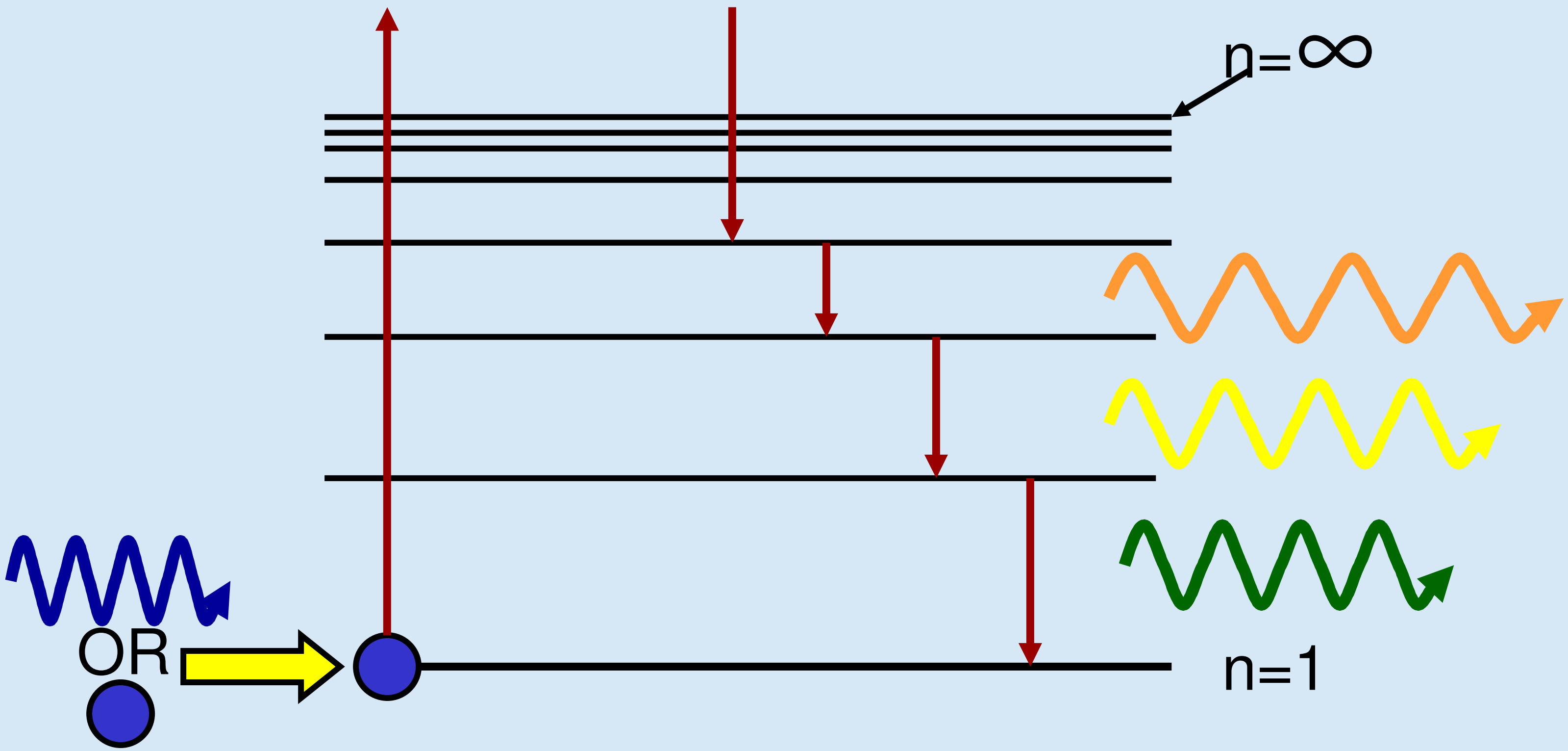
Collisions with electrons/ions/atoms can knock bound electrons into higher energy levels. The energy comes from the kinetic energy of the colliding particle. The electron falls back to lower levels and this energy is radiated away.

Photo-excitation



If a photon with the right energy interacts with an atom or ion, an electron can be moved up to a higher level for a short while, before it falls back down to the ground state.

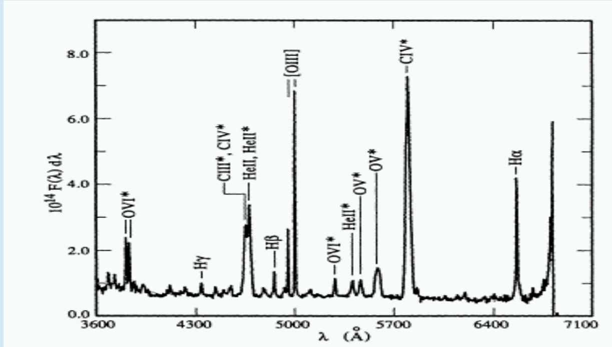
Recombination, following ionization



If a photon or particle with sufficient energy interacts with an atom so that an electron is stripped away, the atom is said to have been ionized. A free electron can recombine with an ion, finishing up into an excited state – it will then cascade down to ground level producing line emission (radiative cascade).

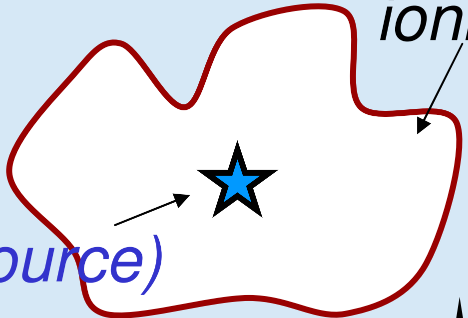
Sources of emission lines

Photoionization ➡ ***Recombination***



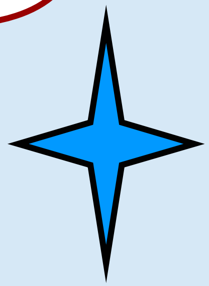
Hot ionized nebulae (e.g. HII regions)

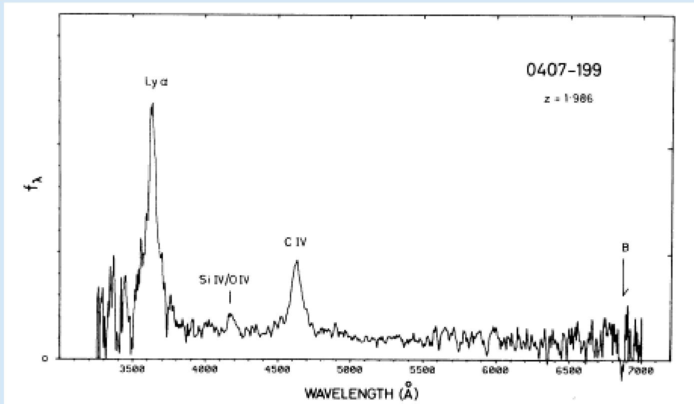
ionized gas



hot star (UV source)

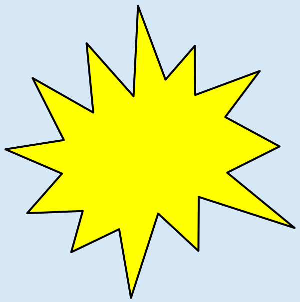
Active galactic nuclei (e.g. quasar)

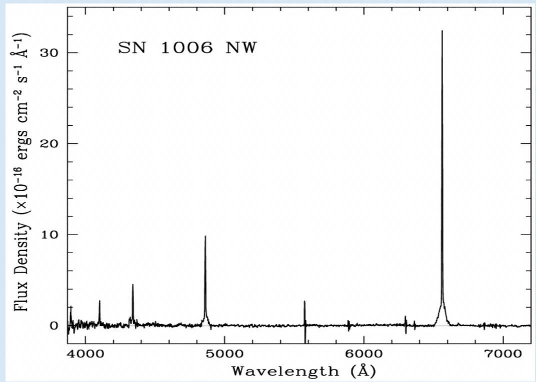




Collisional ionization ➡ ***Recombination***

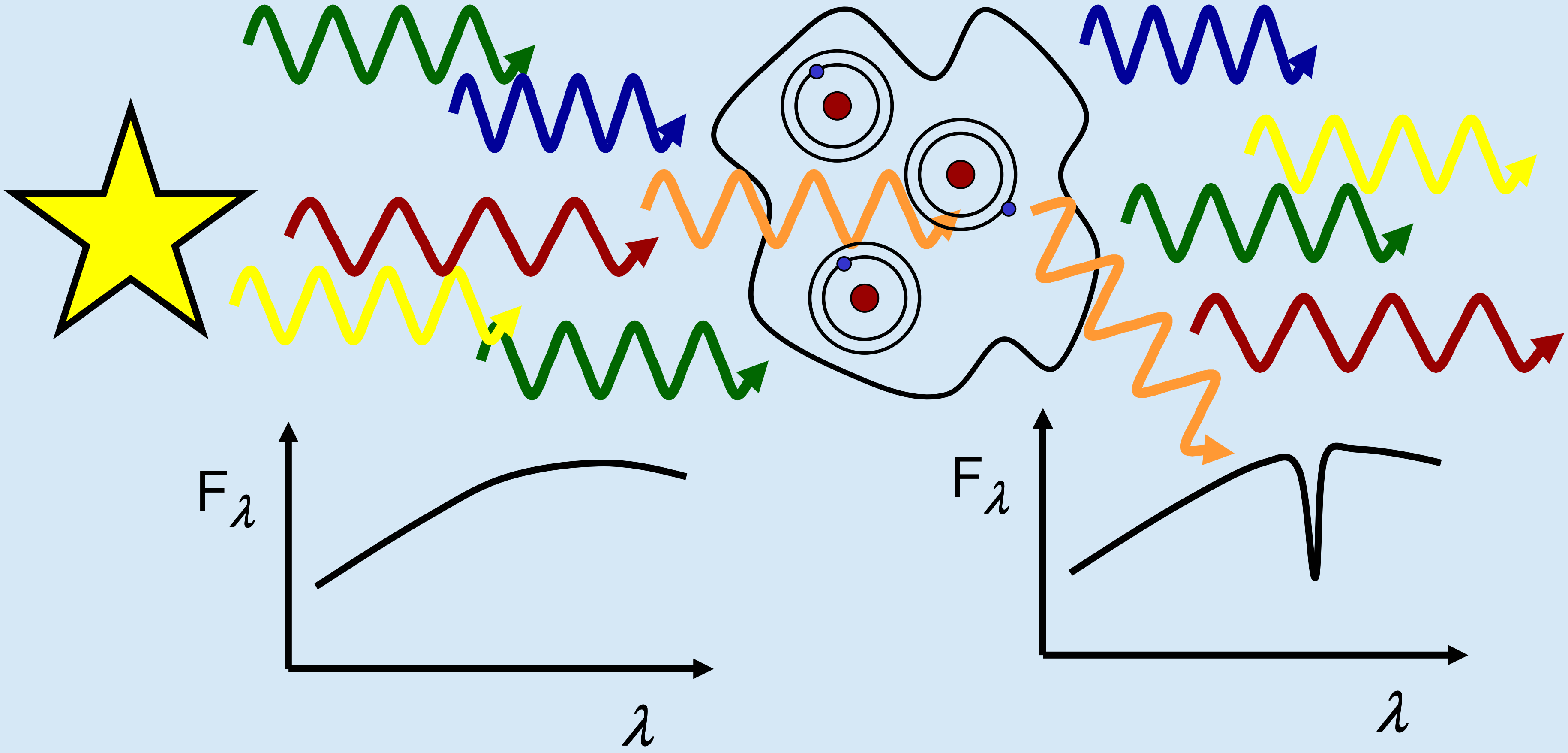
Supernova Remnant





Absorption lines formation

When atoms/ions in a gas are illuminated, they will absorb photons at those wavelengths which will move electrons in the atoms/ions from one level to another.



Absorption lines formation (cont.)

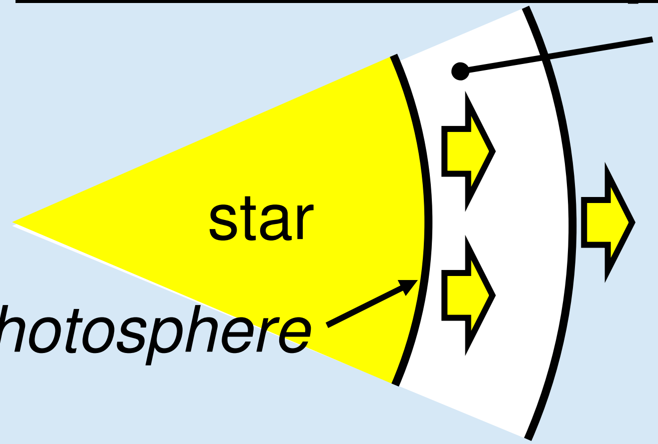
Atoms or ions in a gas will absorb photons whose energy corresponds exactly to the energy that an electron in that atom/ion needs to move into a higher level.

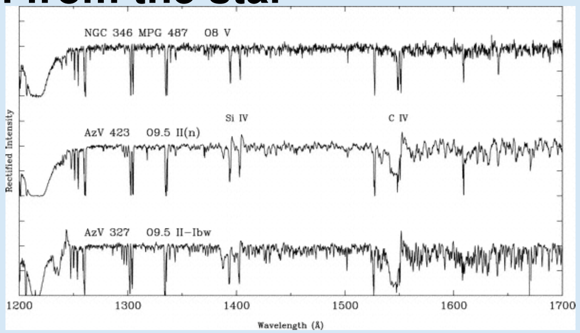
After about 10^{-8} seconds, the electron will fall back down to the most stable state, emitting a photon with an energy corresponding to the difference between the levels, but in a random direction.

So if you look through the gas at a source, you will see very few photons at that energy because these are being absorbed and re-emitted in random directions. This produces an absorption line and explains Kirchoff's 3rd rule.

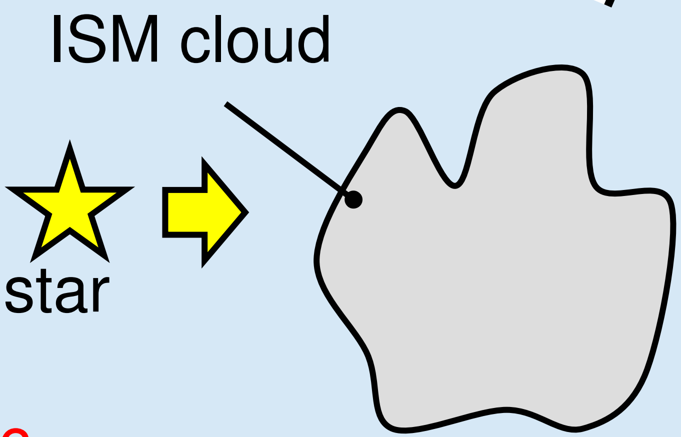
Sources of absorption lines

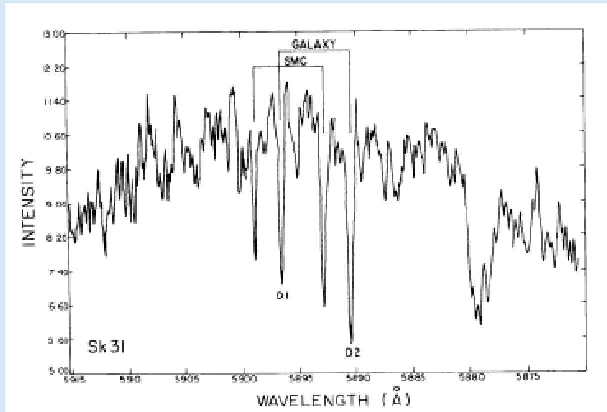
Stellar atmospheres



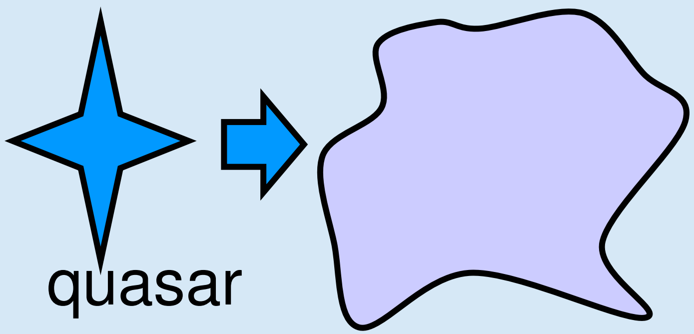


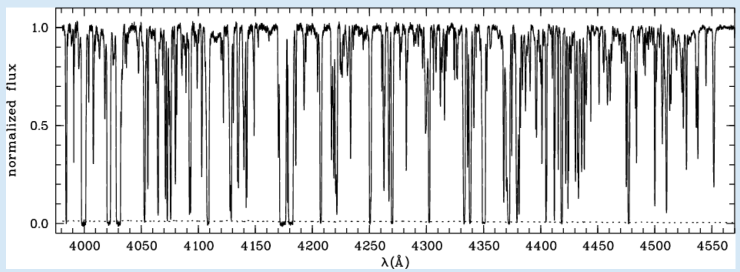
Interstellar gas





Intergalactic Lyα systems of clouds at different redshifts





'Lyα forest'

Stellar classification *(ZG Ch. 13; FK Ch. 19)*

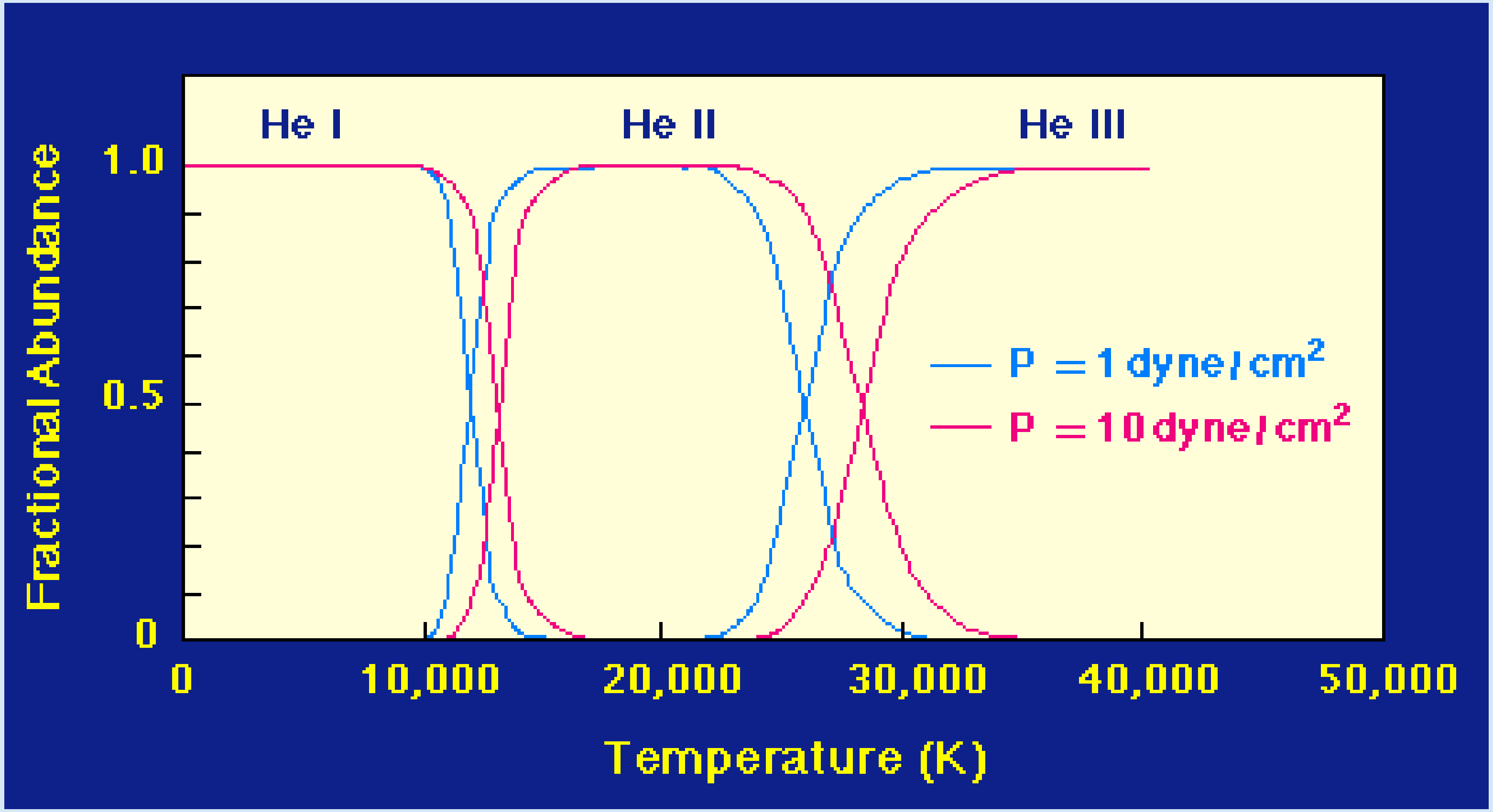
A star is a hot, dense ball of gas. It emits approximately a blackbody spectrum at a single temperature from its photosphere (lowest visible layers).

From Wien's law: Hotter a star is, the bluer. Cool stars look red and intermediate temperature stars appear yellow.

Photosphere is surrounded by thin, warm atmosphere which imprints absorption lines on the continuum.

Which ions do the absorption depends on the temperature of the gas they are in …

Fractional ion abundance (example)



HeI = neutral He atom = $(2p^+ + 2n) + 2e^-$

HeII = He atom once ionised = $(2p^+ + 2n) + e^-$

HeIII = He atom fully ionised = $(2p^+ + 2n)$

$(2p^+ + 2n)$ = α particle

Stellar spectral types (1)

1863: A. Secchi observed prism spectra of ~4,000 stars and divided them into 4 broad spectral classes by common absorption features.

1890: E. Pickering & W. Fleming sorted stars by decreasing hydrogen absorption line strength (spectral type A → B, C, …) … but other lines did not fit into this sequence …

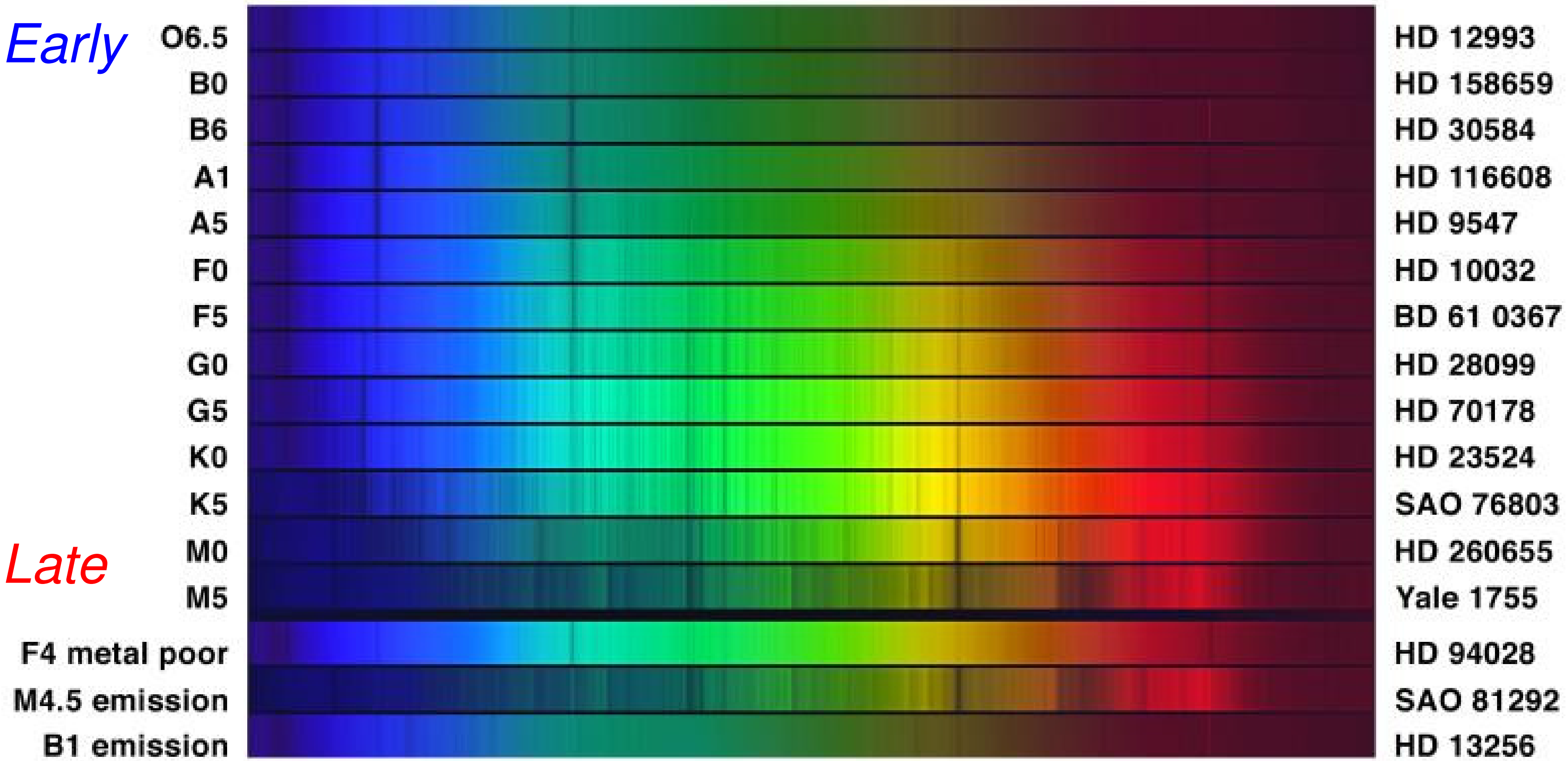
1901: Annie J. Cannon noticed that spectra differ on the basis of stellar surface temperature → re-ordered the ABC types by temperature, refined the classification by dividing each class into 10 subclasses (A1, A2, … A9, …)

➔ From 1910, adopted Harvard (or Henry Draper) spectral classification system, and HD catalogue

Stellar spectral types (2)

O, B, A, F, G, K, M

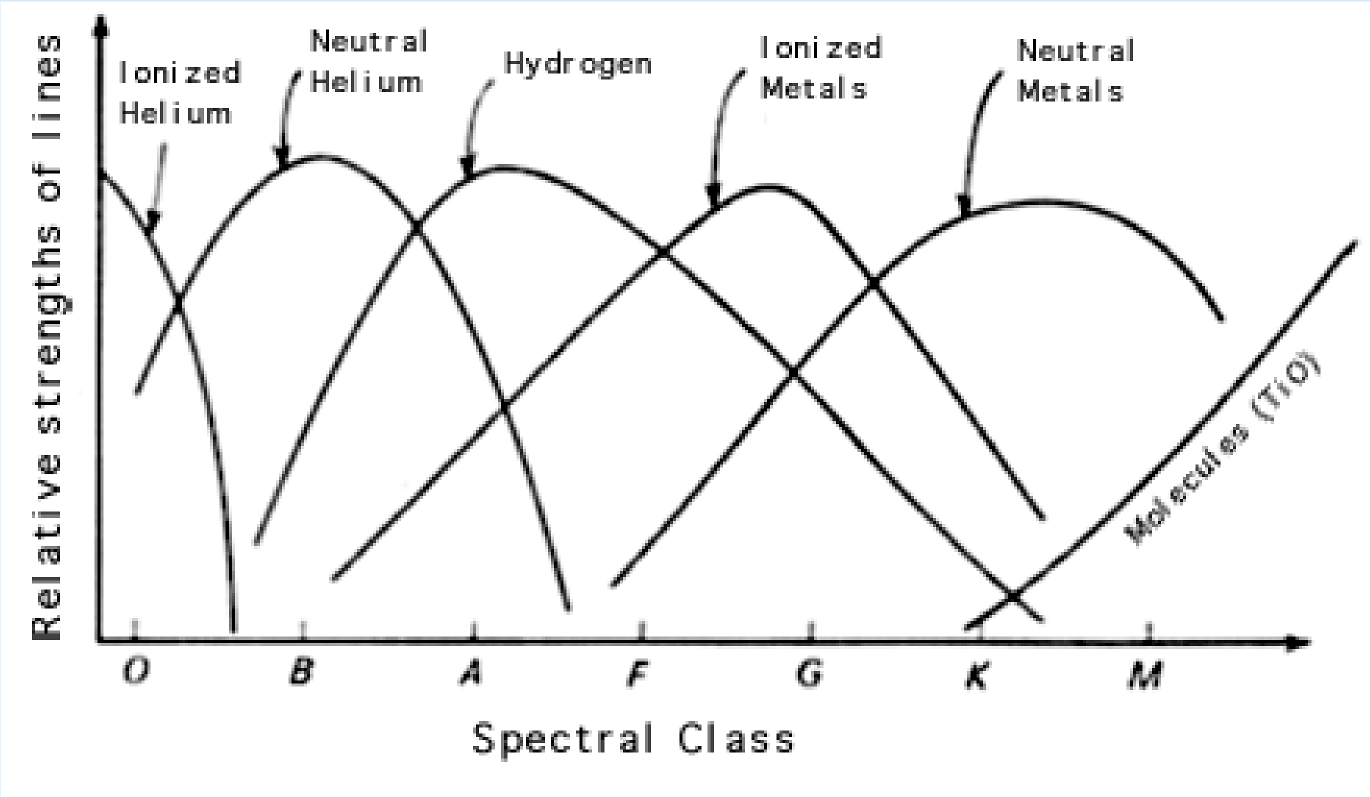
'Early' type ………………… 'Late' type



(evolutionary sequence not implied)

Stellar spectral types (3)

Which ions do the absorption depends on the temperature of the gas they are in …



Stellar spectral types (4)

Spectral class	Colour	Surface temp (K)	Main lines	Example
O	Blue-violet	30000-50000	He II	Naos
B	Blue-white	11000-30000	He I	Rigel, Spica
A	White	7500-11000	H, Fe II, Si II, Mg II	Sirius, Vega
F	Yellow-white	6000-7500	Ca II	Canopus, Procyon
G	Yellow	5000-6000	Ca II, Fe I, CH	Sun, Capella
K	Orange	4000-5000	Ca II, Fe I	Arcturus, Aldebaran
M	Red-orange	<4000	Fe I, TiO	Betelgeuse, Antares

Stellar spectral types (5)

1943: Morgan & Keenan added luminosity as a second classification parameter.

Luminosity classes are designated by the Roman numerals I thru V, in order of decreasing luminosity:

- Ia = Most luminous supergiants
- Ib = Supergiants
- II = Luminous giants
- III = Giants
- IV = Subgiants
- V = Main sequence stars (dwarfs)

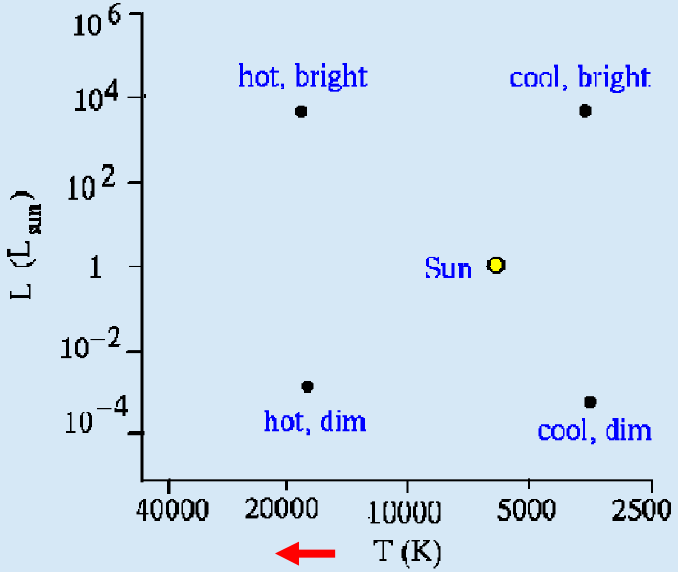
The Sun is a G2V star and has an effective (surface) temperature of ~6000 K.

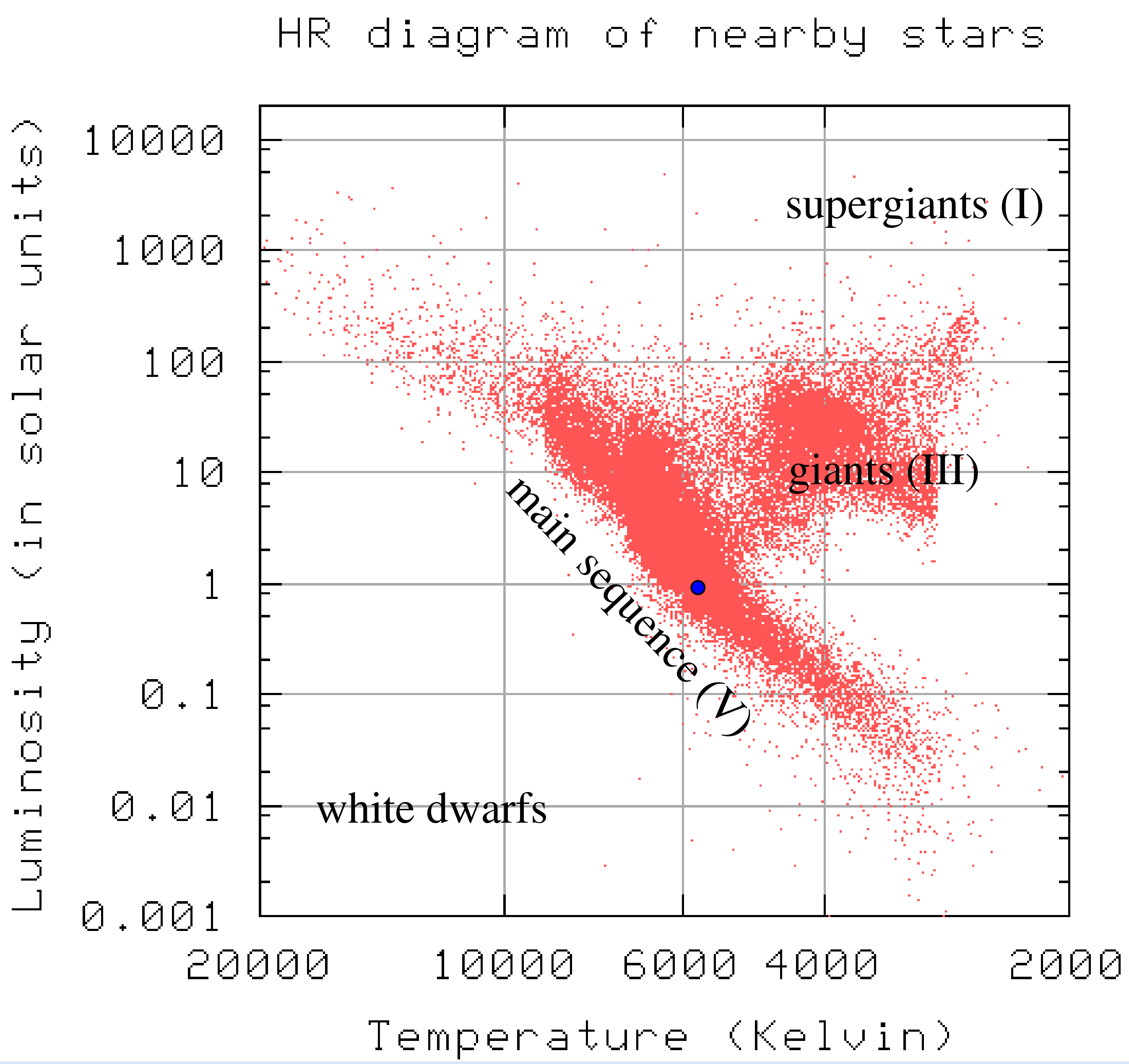
Hertzsprung-Russell diagrams

1911-1913: Hertzsprung and Russell independently plotted stellar luminosity vs temperature (or spectral type).

Stars populate the diagram preferentially in certain regions. This may be partially understood in terms of luminosity of an object emitting thermal radiation:

$$L \sim R^2 T^4$$




HR diagram of nearby stars
Luminosity (in solar units)
10000
1000
100
10
1
0.1
0.01
0.001
20000
10000
6000
4000
2000
Temperature (Kelvin)
supergiants (I)
giants (III)
main sequence (V)
white dwarfs