T Tauri stars

Low-mass <u>pre</u> main sequence stars

Transition objects between earlier stages of collapse, ~~[illegible]~~ still shrouded in dust & only visible in IR, and main sequence stars

Unstable environment : mass loss ("P Cygni profiles"), rapid variations in luminosity, etc

These stars are found to be on the pre-mainsequence evolutionary tracks when T, L are measured.

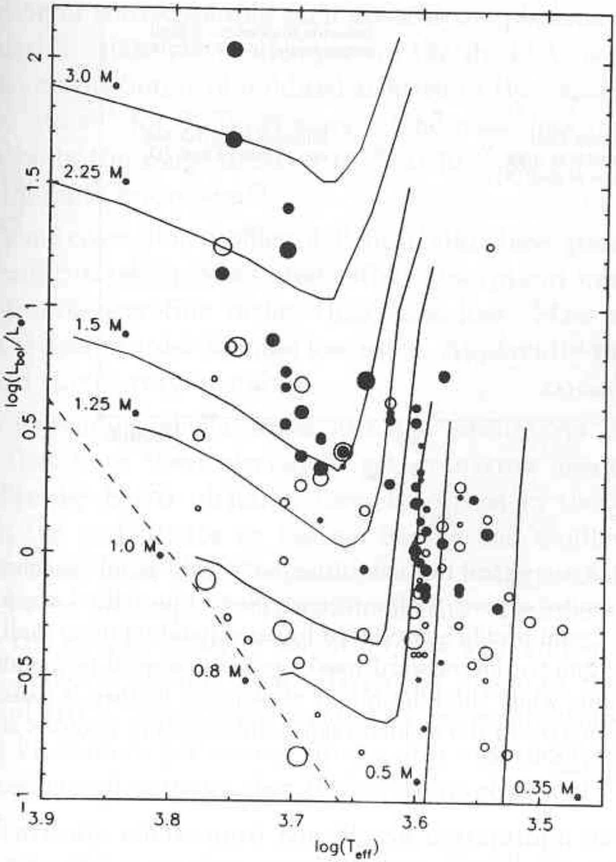


Figure 12.11 The positions of T Tauri stars on the H–R diagram. The sizes of the circles indicate the rate of rotation. Stars with strong emission lines are indicated by filled circles and weak emission line stars are represented by open circles. Theoretical pre-main-sequence evolutionary tracks are also included. (Figure from Bertout, *Annu. Rev. Astron. Astrophys.*, *27*, 351, 1989. Reproduced with permission from the *Annual Review of Astronomy and Astrophysics*, Volume 27, ©1989 by Annual Reviews Inc.)

Early stages of collapse of protostar

Jeans mass : minimum mass for grav. collapse

$$M_J \propto T^{3/2} \rho^{-1/2}$$

Early collapse is isothermal as cloud density low & it is optically thin

$\boxed{Q.}$ What happens to the Jeans mass as the isothermal collapse proceeds?

$\rho \uparrow$ so $M_J \downarrow$

Smaller regions (smaller masses) become gravitationally unstable

$\longrightarrow$ FRAGMENTATION

This is why stars (particularly young ones) are found in clusters

Youngest clusters are known as O-B associations

$\boxed{Q}$ It is likely that stars of all masses are formed, and in fact low-mass stars are more common. Why are they called O-B ass^{ns} ?

Many young star clusters disperse as they become older, but dense clusters or those that live away from the disk most of the time can survive

OB associations - groups of very young high mass stars (some lower mass stars too)

- not gravitationally bound; expanding & will dissolve into field.

How is it possible for a gravitationally bound cloud to give birth to a gravitationally unbound association?

Evolution after the main sequence

Once a star arrives on the main sequence, it is burning $H \rightarrow He$ in its core.

Mean molecular weight in core is slowly increasing

Recall ideal gas law

$$P = \frac{\rho k T}{\underbrace{\mu}_{\text{mean molecular weight}} m_H}$$

Q If ρ, T stay unchanged, what happens to the pressure? How will the star's structure adjust?

If pressure in core decreases, it will not be able to support overlying layers of the star.

So, the core will be compressed.

Gravitational potential energy released

$\boxed{\text{Virial theorem}}$: half energy is radiated away, half goes into increasing temperature.

Nuclear reaction rate for p-p chain goes as $\rho X^2 T_6^{\ 4}$ (near 1.5×10^7 K)

X is mass fraction of hydrogen

T_6 is temp in units of 10^6 K.

Mass fraction of hydrogen is decreasing as $H \rightarrow He$

BUT $\rho \uparrow$ and $T \uparrow$ and these terms win

Reaction rate goes up and so luminosity of star slowly increases

$\rightarrow$ Star leaves the Zero Age Main Sequence

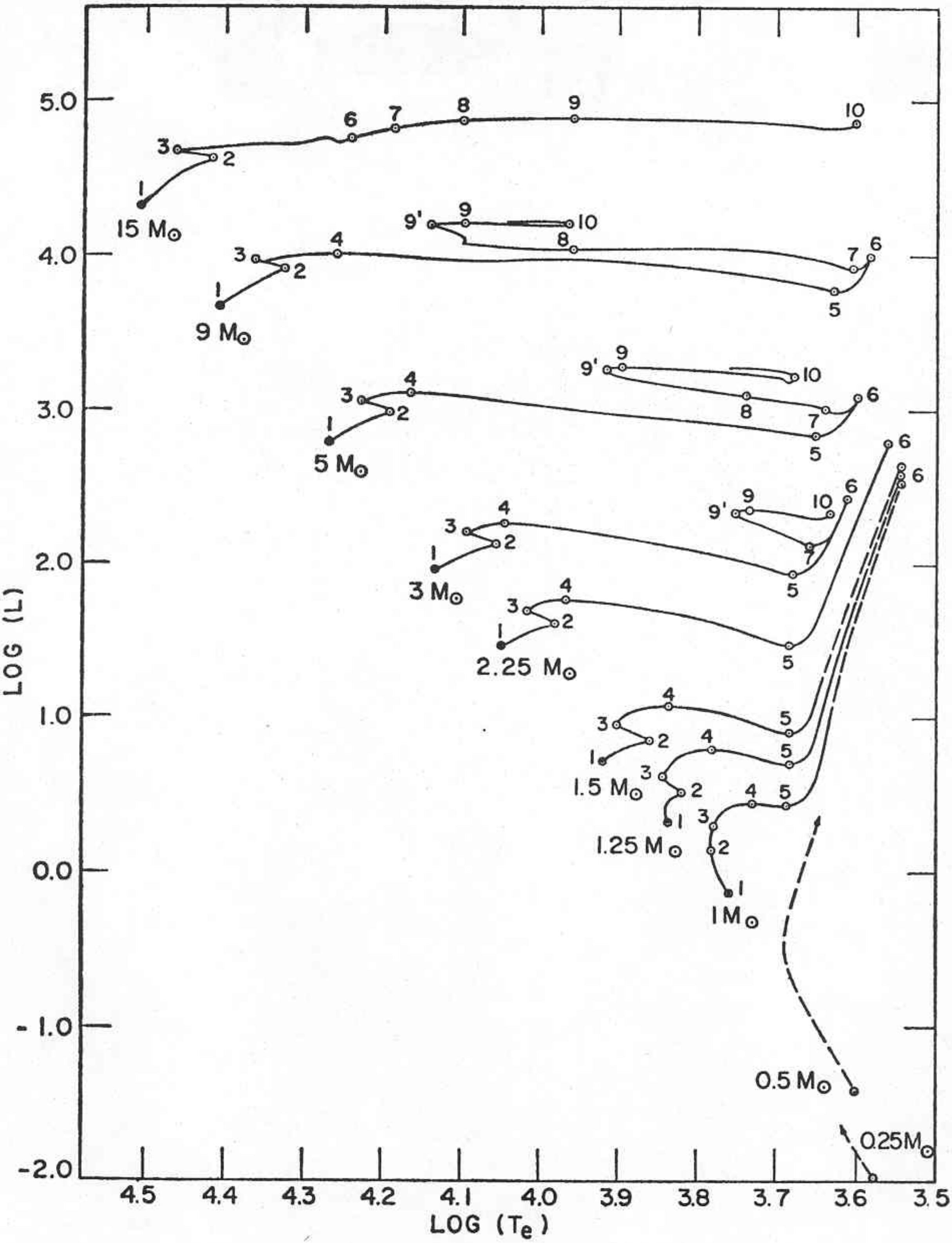


Figure 13.1 Main-sequence and post-main-sequence evolutionary tracks of stars with an initial composition of $X = 0.708$, $Y = 0.272$, and $Z = 0.020$. (Figure from Iben, *Annu. Rev. Astron. Astrophys.*, *5*, 571, 1967. Reproduced with permission from the *Annual Review of Astronomy and Astrophysics*, Volume 5, ©1967 by Annual Reviews Inc.)

Why stars leave the main sequence (properly)

Stable main-sequence star is in equilibrium:

inward force of gravity

balanced by

outward force of gas & radiation pressure

(caused by fusion)

But it's core, where fusion is going on, is slowly being turned from hydrogen to helium.

- What happens when there is no hydrogen left in the core?
- Can we see the helium increase in the star's spectrum?

Fairly quickly, in a star the mass of the Sun, this gravitational heating will heat the shell of hydrogen just outside the core & start fusion there.

called **hydrogen shell burning.**

A star's luminosity is determined by its surface temperature & surface area. At point 3 in the diagram the star's luminosity is slightly higher but the temperature is almost unchanged. What does this imply about the star's diameter?

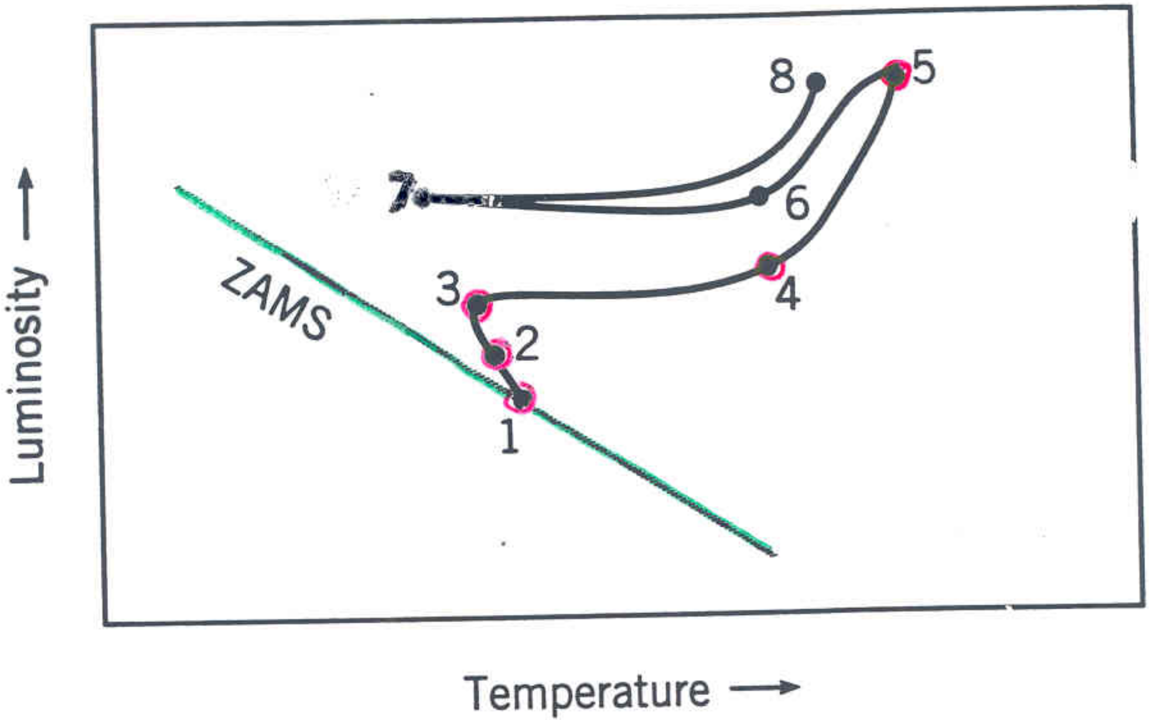


FIGURE 18–3. A schematic diagram of the evolutionary track of a star of one solar mass in the Hertzsprung-Russell diagram. The events occurring at each numbered point are discussed in the text.

- Most of photons generated in hydrogen-burning shell do not easily escape the star, but are absorbed by outer layers & cause them to expand

core contracts H burning in shell

outer layers expand - energy from shell

- So as envelope expands, it cools

(just as gases expanding out of aerosol can cool)

- So temperature of star's surface decreases (4)

- Then whole envelope becomes convective, energy can get out, star gets more luminous & becomes a red giant (5)

Evolution during red giant phase is FAST compared to main sequence lifetime ...... so red giants are relatively rare.

<u>Question</u> : there are several well-known bright red giants in the night sky — how can this be if red giants are rare ?

Core continues to contract and heat until it reaches a temperature of 10^8 K — ~~xxxx~~ ten times its main sequence temperature

This is hot enough for helium nuclei to fuse to make carbon

$${}^4_2He + {}^4_2He \longrightarrow \left({}^8_4Be\right) \qquad (1)$$

unstable

very soon after

$${}^8_4Be + {}^4_2He \longrightarrow {}^{12}_6C + \gamma \qquad (2)$$

TRIPLE ALPHA reaction

(5)

Need high enough ρ, T for reaction 2 to happen before Be decays

Helium ignition can happen so quickly that it is explosive (HELIUM FLASH)

if the core is so dense that it becomes <u>degenerate</u>

In this case, increasing temperature does not increase pressure as in an ideal gas (normally in a star)

So star does not expand (as usual)

We call this a <u>thermal runaway</u>

<u>Question</u> What is likely to happen when all the core He is burnt to carbon?

Low-mass stars (cont)

Horizontal branch : core He burning

shell H burning

Asymptotic giant branch (AGB)

: C-O core

: shell He burning

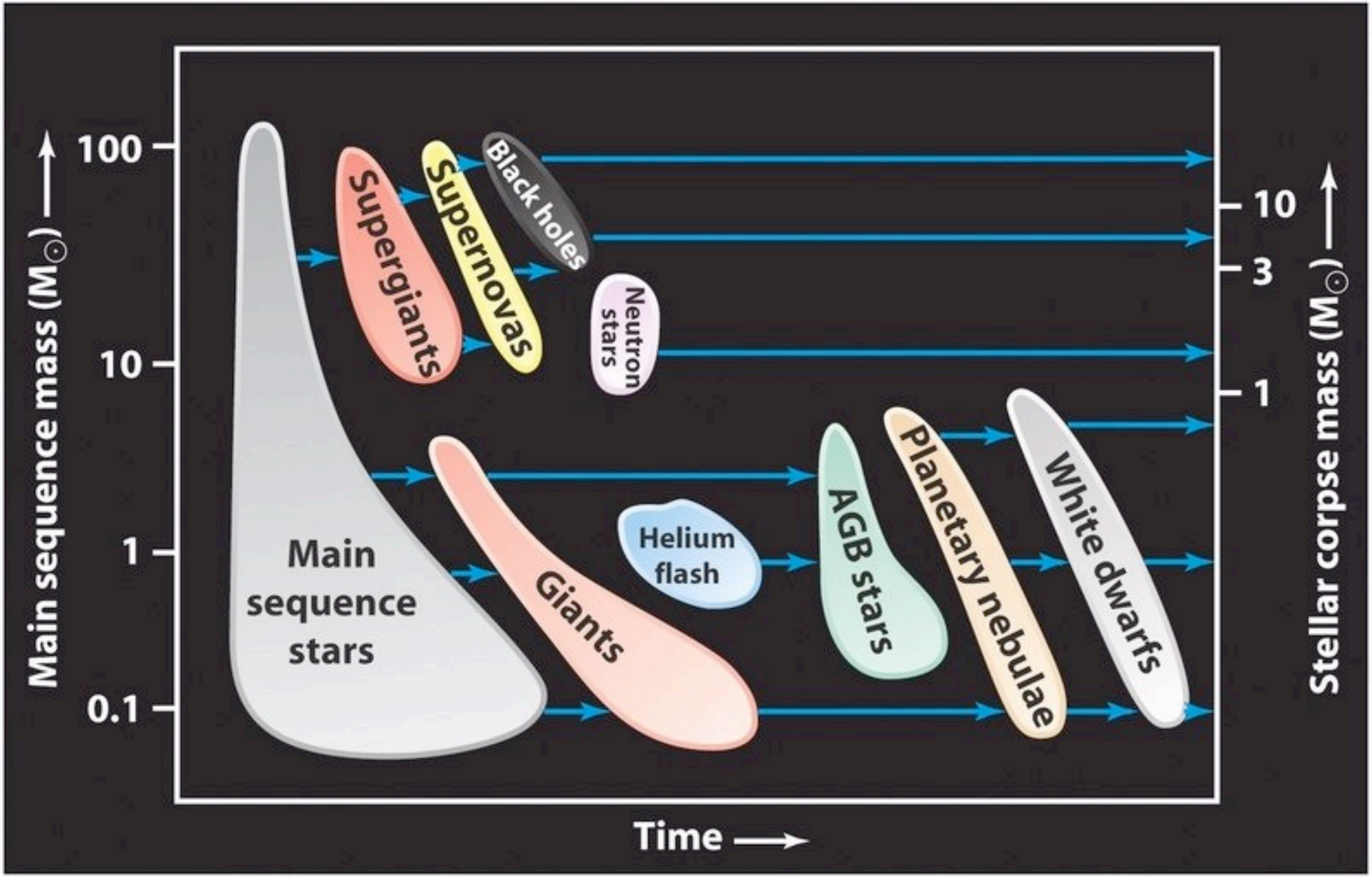
: shell H burning

- Stars pulsate (long-period variables) at & near tip of AGB

- Star loses remaining H envelope, puffs off shells; seen in IR & radio (OH/IR stars) and as "planetary nebulae".

- Central star has little H left; never gets hot enough to ignite C & O.

Becomes a white dwarf.


Main sequence mass ($M_\odot$)
100
10
1
0.1
Main sequence stars
Supergiants
Supernovas
Black holes
Neutron stars
Giants
Helium flash
AGB stars
Planetary nebulae
White dwarfs
Stellar corpse mass ($M_\odot$)
10
3
1
Time

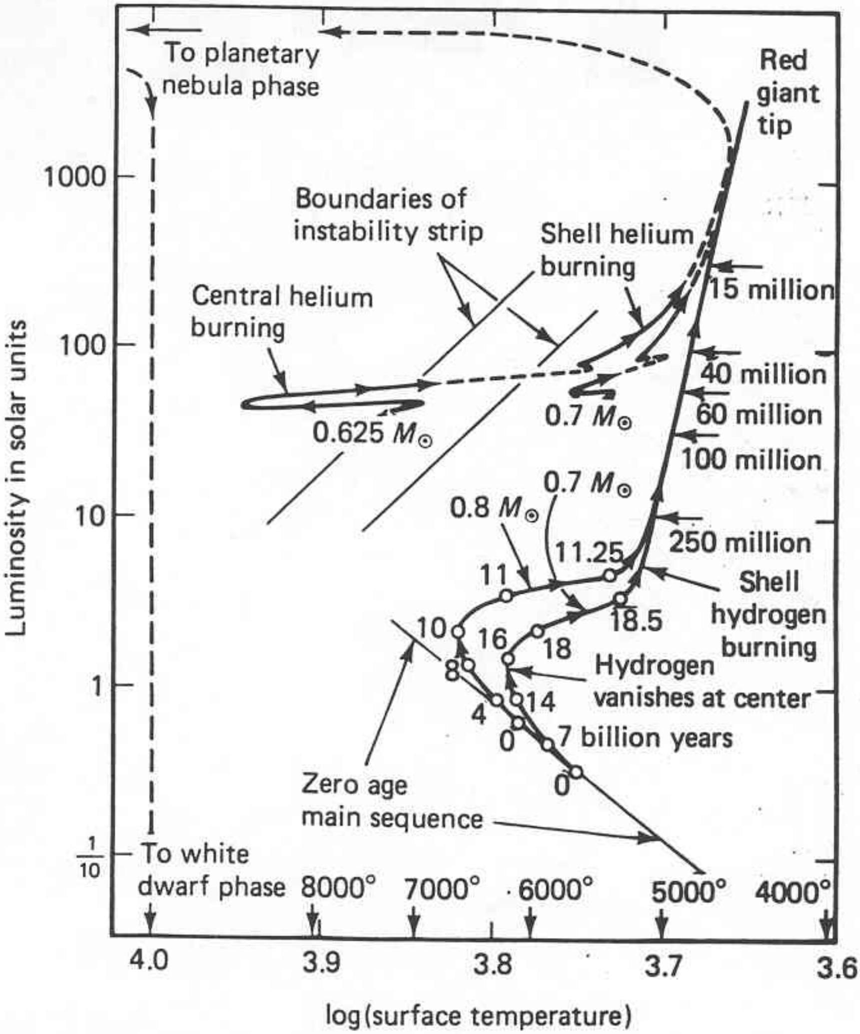


Fig. 14.11. Evolutionary tracks for population II stars with 0.7 and 0.8 $M_\odot$. A helium abundance of $Y = 0.30$ and a heavy element abundance of $Z = 10^{-3}$ was used. On the main sequence and subgiant branches evolution times since arrival on the zero age main sequence are given in billions of years. On the red giant branch evolution times from one arrow to the next are given in millions of years. Also indicated is the instability strip, the T_{eff}, L domain in which stars start to pulsate (see Chapter 17). From the tip of the asymptotic giant branch the stars probably evolve through the planetary nebula stage to become white dwarfs as indicated by the long dashed lines. From Iben (1971).

White dwarfs

Final stage of evolution of low-mass stars (central star of planetary nebula becomes white dwarf)

~ size of Earth

Q Why do you think the error bars are so large in the top part of the diagram?

Types of white dwarfs

temp can range from 5000 K → 80,000 K

DA – most common
– broad H lines only in spectrum

DB – no H lines, only He

DC – no lines at all!

— almost all C & O —

White dwarf cooling

There are no nuclear reactions in a white dwarf — it is a ball of inert C & O which may have a small H/He atmosphere

$\boxed{Q}$ What is the source of its luminosity ?

→ Thermal energy $U = \frac{3}{2} N k T_c$

interior temp

We can derive luminosity from

$$L = -\frac{\partial U}{\partial t}$$

and eq'ns of stellar structure

Find that

$$L_{wd} = L_0 \left(1 + \frac{5}{2} \frac{t}{\tau_0} \right)^{-7/5}$$

τ_0 is cooling timescale @ T_0 ~ 2×10^8 yr

We can use this to estimate the age of clusters and of the stellar disk (and compare with stellar evolution ages)

As nuclei cool they eventually settle into a solid :

'crystallization'

Electron degeneracy

- No 2 electrons can occupy exactly the same quantum state (Pauli exclusion principle)

- Uncertainty principle:

  $\Delta p \, \Delta x \gtrsim \hbar$

  (can be thought of as defining available quantum states)

A classical (ideal) gas will have no motion at $T = 0$

A degenerate electron gas does ....

Very high density e^- gas :

$\Rightarrow$ Δx between neighboring electrons very small

$\Rightarrow$ Δp must be correspondingly large

- Electrons must have momenta that differ by $\geqslant \hbar/\Delta x$ or Pauli exclusion principle is violated (strictly, can have 2 because of spin)

- So some electrons have very high momenta

- This leads to a high pressure which is independent of ~~[illegible]~~ temperature in the case of complete degeneracy.

$P \propto \rho^{5/3}$ (non-relativistic e^- degeneracy)

& $P \propto \rho T$ (ideal gas)

How dense do we need to be?

White dwarfs have density $\sim 10^6 \, g/cm^3$

at 10^7 K, degeneracy pressure is

$\sim 100 \times$ thermal pressure!

Mass-radius relation

The more massive a white dwarf, the smaller it is:

Hydrostatic $\equiv^m$ $\frac{dP}{dr} = -\rho g_r$

one shell approx. $\frac{P_c}{R} = \frac{M}{\frac{4}{3}\pi R^3} \frac{GM}{R^2}$

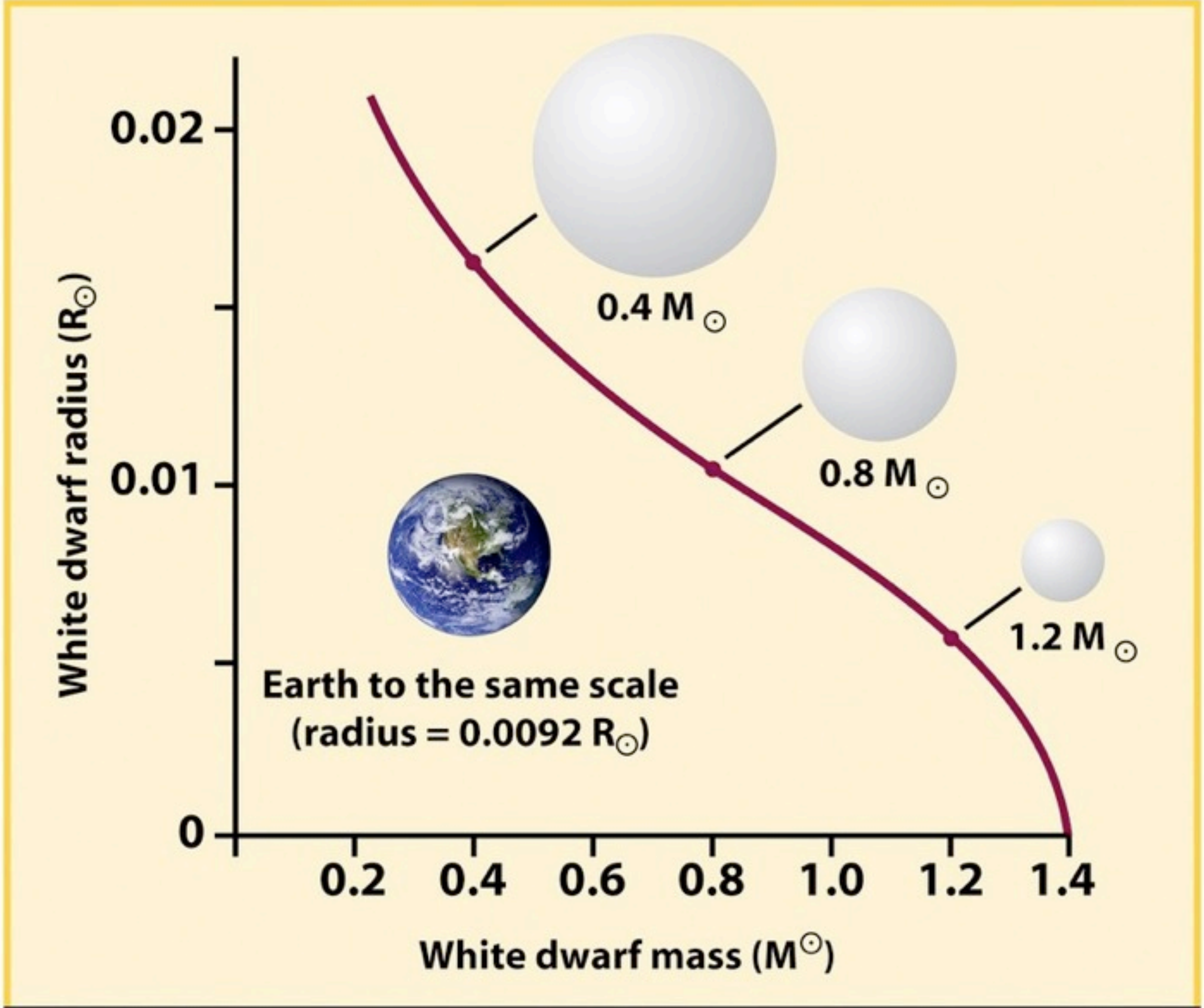
$$P_c \simeq \frac{GM^2}{R^4}$$

Now $P \propto \rho^{5/3} \propto \left(\frac{m}{R^3}\right)^{5/3}$

So $\left(\frac{m}{R^3}\right)^{5/3} \propto \frac{gm^2}{R^4}$

$$\frac{m^{5/3}}{R^5} \propto \frac{gm^2}{R^4}$$

$\Rightarrow \quad R \propto m^{-1/3}$



0.02
0.01
0
White dwarf radius ($R_\odot$)
0.4 $M_\odot$
0.8 $M_\odot$
1.2 $M_\odot$
Earth to the same scale
(radius = 0.0092 $R_\odot$)
0.2
0.4
0.6
0.8
1.0
1.2
1.4
White dwarf mass ($M^\odot$)

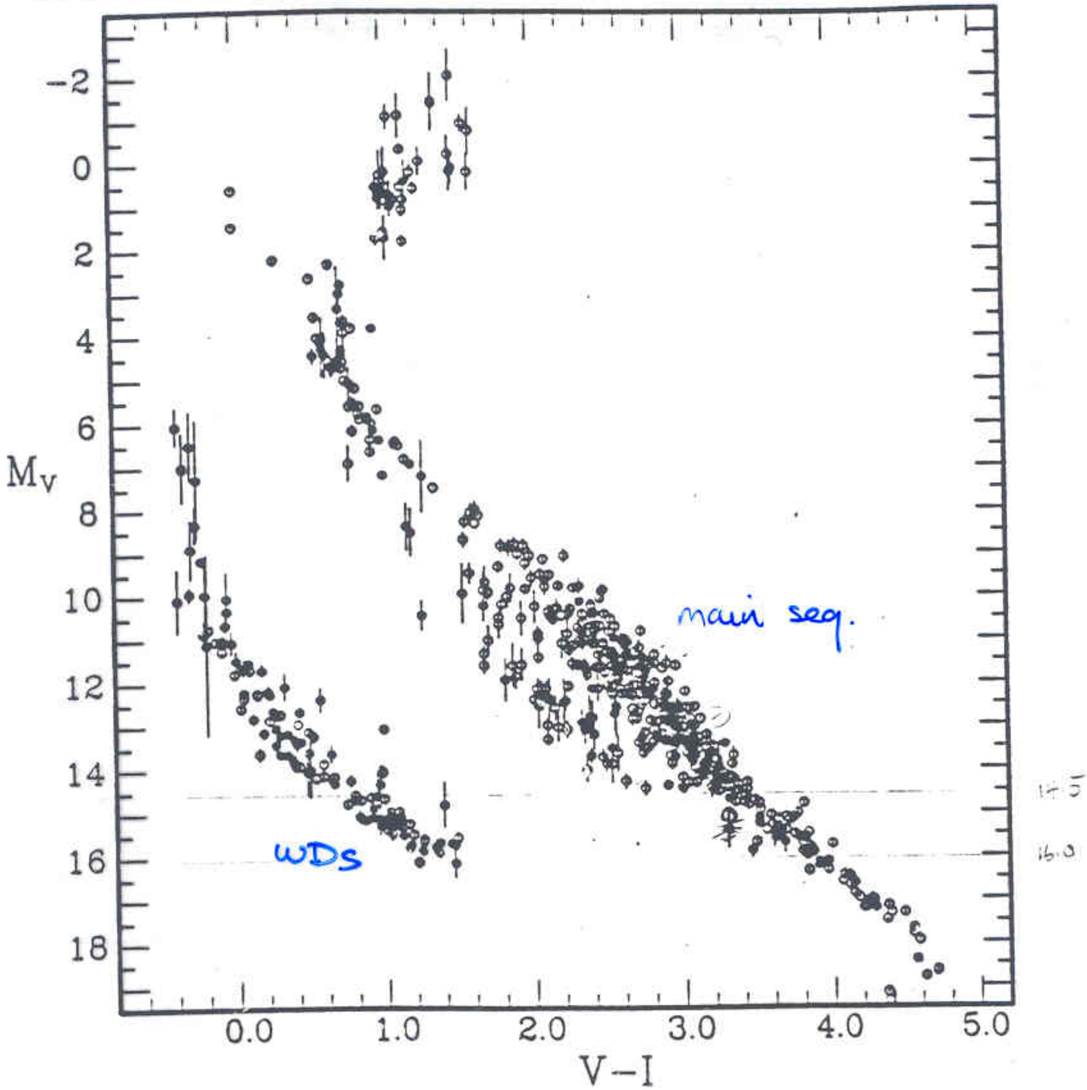


Fig. 1.1. The M_V vs V−I color-magnitude diagram for a selection of 680 stars with (*a*) published USNO parallaxes (open circles; see text) or (*b*) unpublished but completed or nearly completed USNO parallaxes (filled circles).

Fig. 1.1 shows the M_V vs V−I color-magnitude diagram for a selection of 510 stars with published USNO parallaxes along with 170 of the completed or nearly completed stars. The stars with published parallaxes in the figure were selected for meeting the following criteria: formal mean errors in M_V of a) $\leq \pm 0.20$ for the dwarfs, b) $\leq \pm 0.40$ for the subdwarfs, and c) $\leq \pm 0.15$ for the degenerates. (Giants and subgiants are shown, irrespective of the parallax quality, due to the small numbers of such stars.) All points are plotted with $\pm$ 1 m.e. error bars but, due to the scale of the diagram, they often do not show.

There is a limit to the mass that e^- degeneracy pressure can support against gravity - WHY ?

Chandrasekhar limit

We can estimate the speed of electrons in a white dwarf like Sirius B :
(see C&O for calculation)

$$v \simeq 1.1 \times 10^{10} \text{ cm/s}$$

ie $\frac{1}{3}$ speed of light !

If star is more massive, density $\uparrow$ and v $\uparrow$

Finite speed of light imposes limit on mass of a white dwarf

Chandrasekhar limit $M = 1.44\ M_\odot$

$\boxed{Q}$ Imagine a white dwarf star close to the Chandrasekhar limit in a close binary system, accreting mass from its companion. What might happen?

This is one theory for the origin of a Type I supernova.