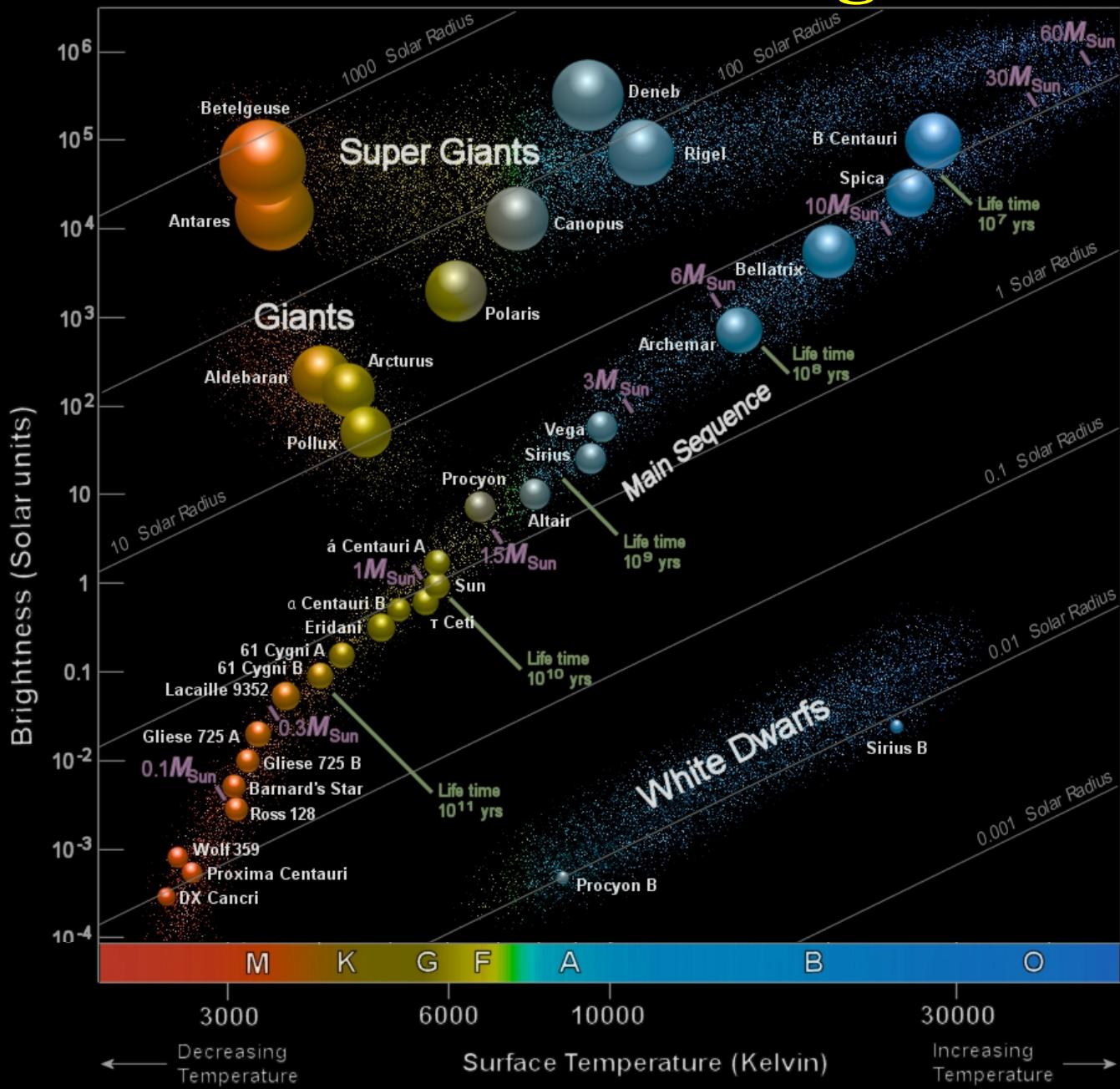


Ay 1 – Lecture 9

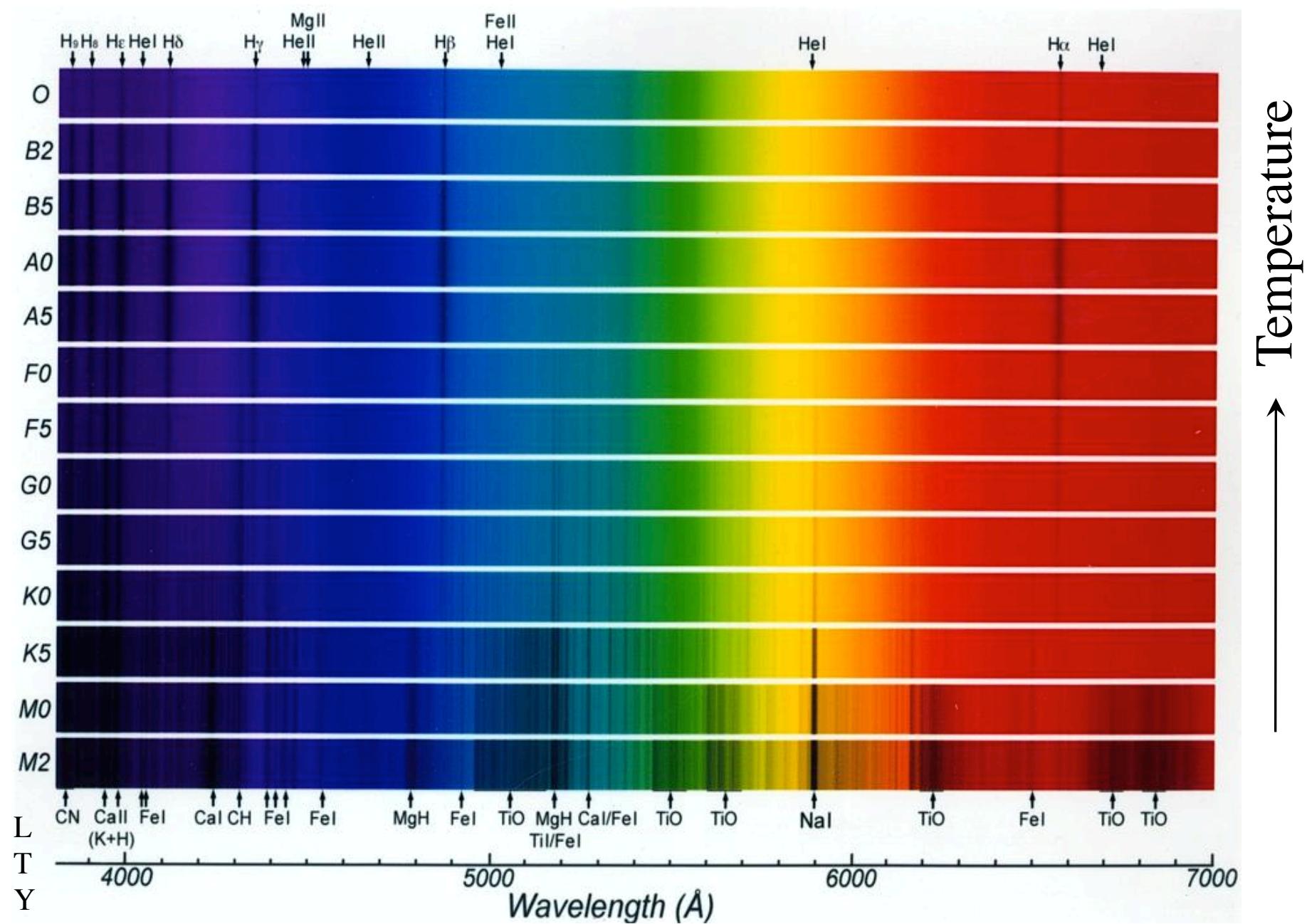
HR Diagram, Star Clusters, and Stellar Evolution

M7 ESO

9.1 The HR Diagram

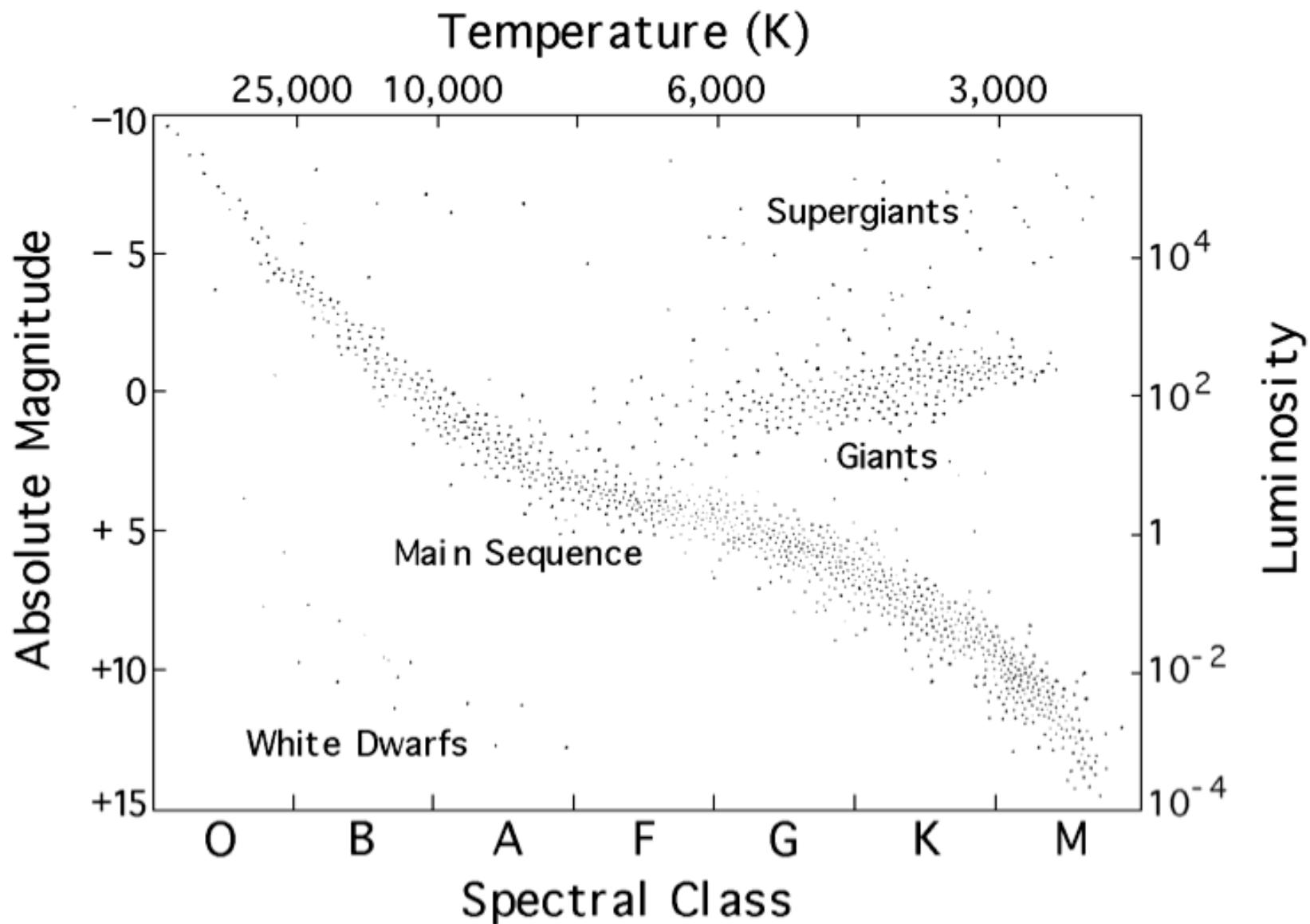


Stellar Spectral Types



The Hertzsprung-Russel (HR) Diagram

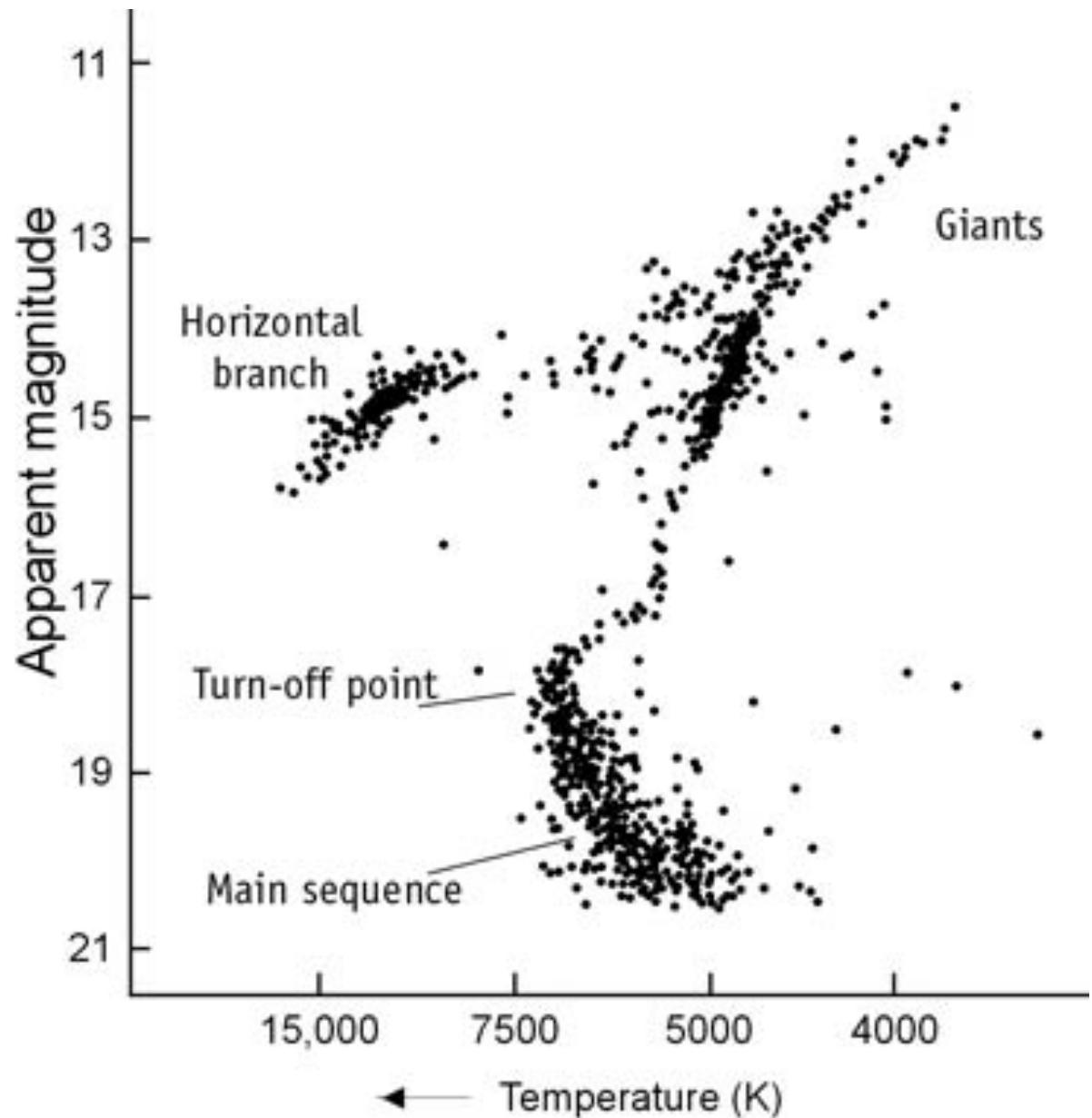
It is a plot of stellar luminosity vs. temperature (~ spectral type)



The Hertzsprung-Russel (HR) Diagram

Stars form *distinctive sequences* in the HR diagram

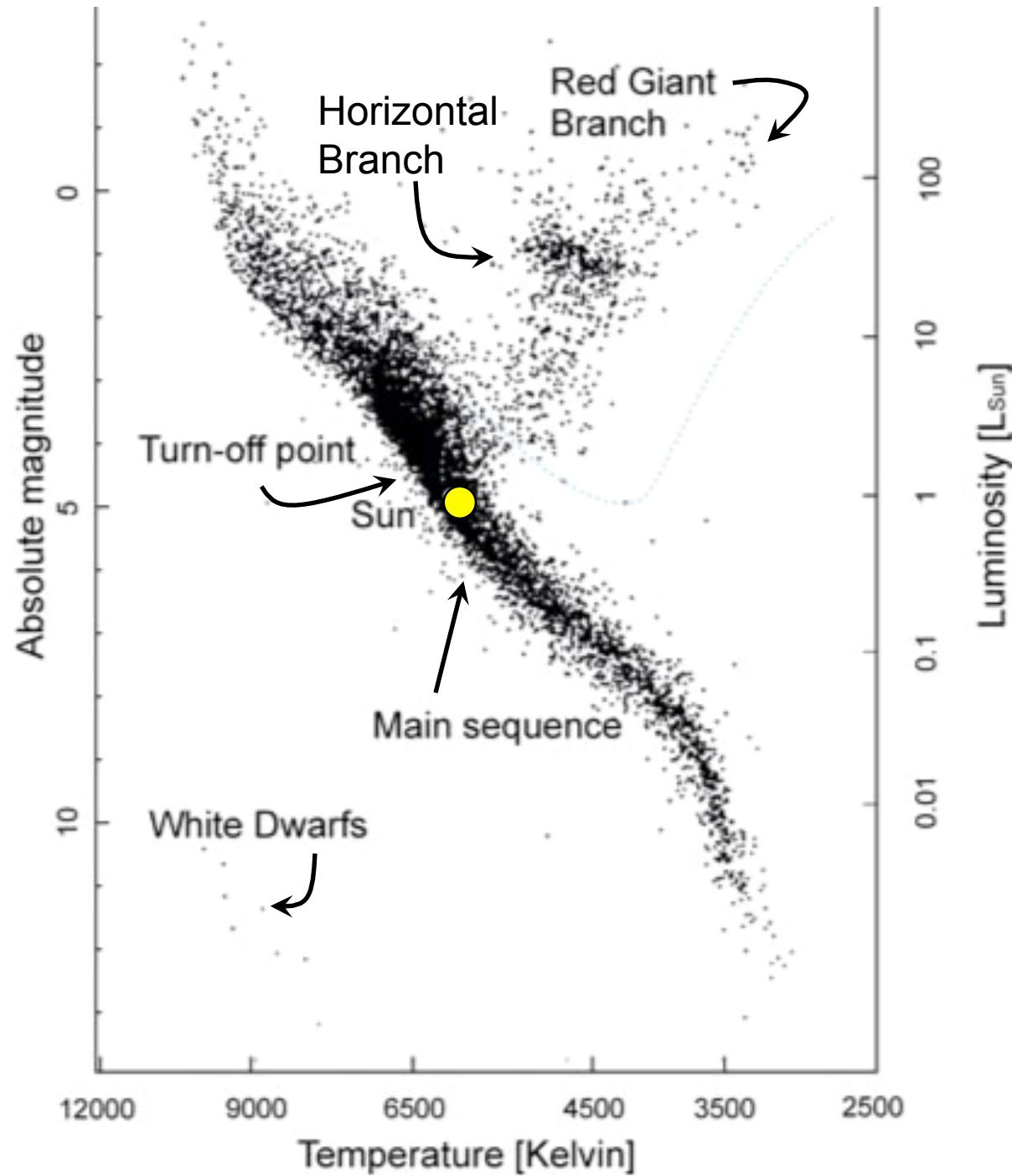
This can be understood in the context of the stellar evolution models, and used to test them



The Principal Sequences

They correspond to different stages of the stellar evolution

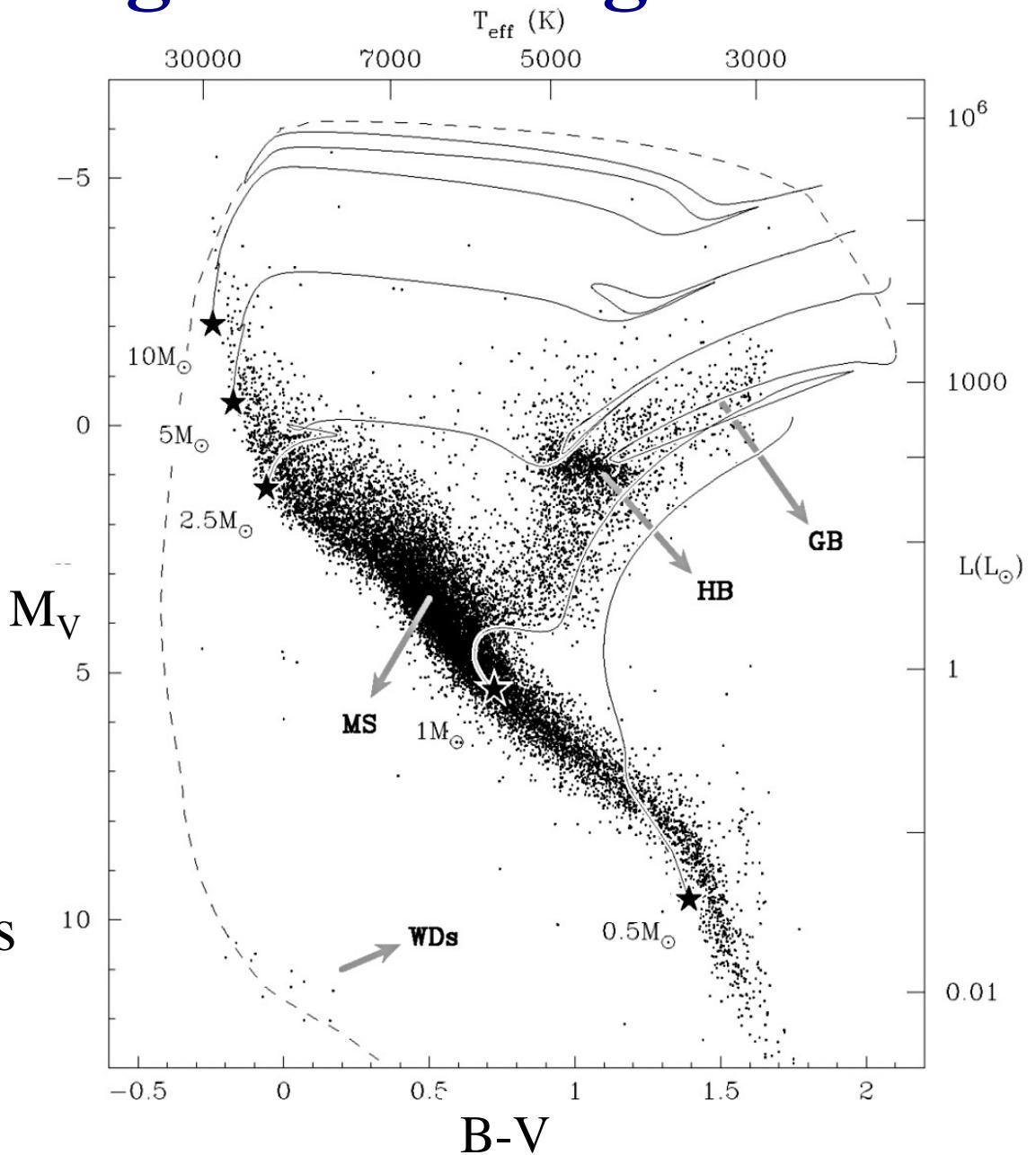
Note: to get the absolute magnitudes (or luminosities), you must have *distances*



The Color-Magnitude Diagram

Generally, photometric colors are used on the X axis, as a measure of temperature, e.g., $(B-V)$ = difference in magnitudes in two different filters (B and V)

Need models of stellar atmospheres to convert these measured quantities into temperatures and bolometric luminosities



9.2 Star Clusters: Probes of Stellar Evolution

NGC 3603 ESO

Star Clusters

Open (or Disk):

$N_\star \sim 10^2 - 10^3$
Ages $\sim 10^7 - 10^9$ yr



Pleiades

Globular:

$N_\star \sim 10^4 - 10^7$
Ages $\sim 10 - 13$ Gyr



M2

- Great “laboratories” for stellar evolution and dynamics
- Dynamical and evolutionary time scales $<$ or \ll Galaxy’s age, and a broad range of evolutionary states is present



M25 ESO



M13 HST

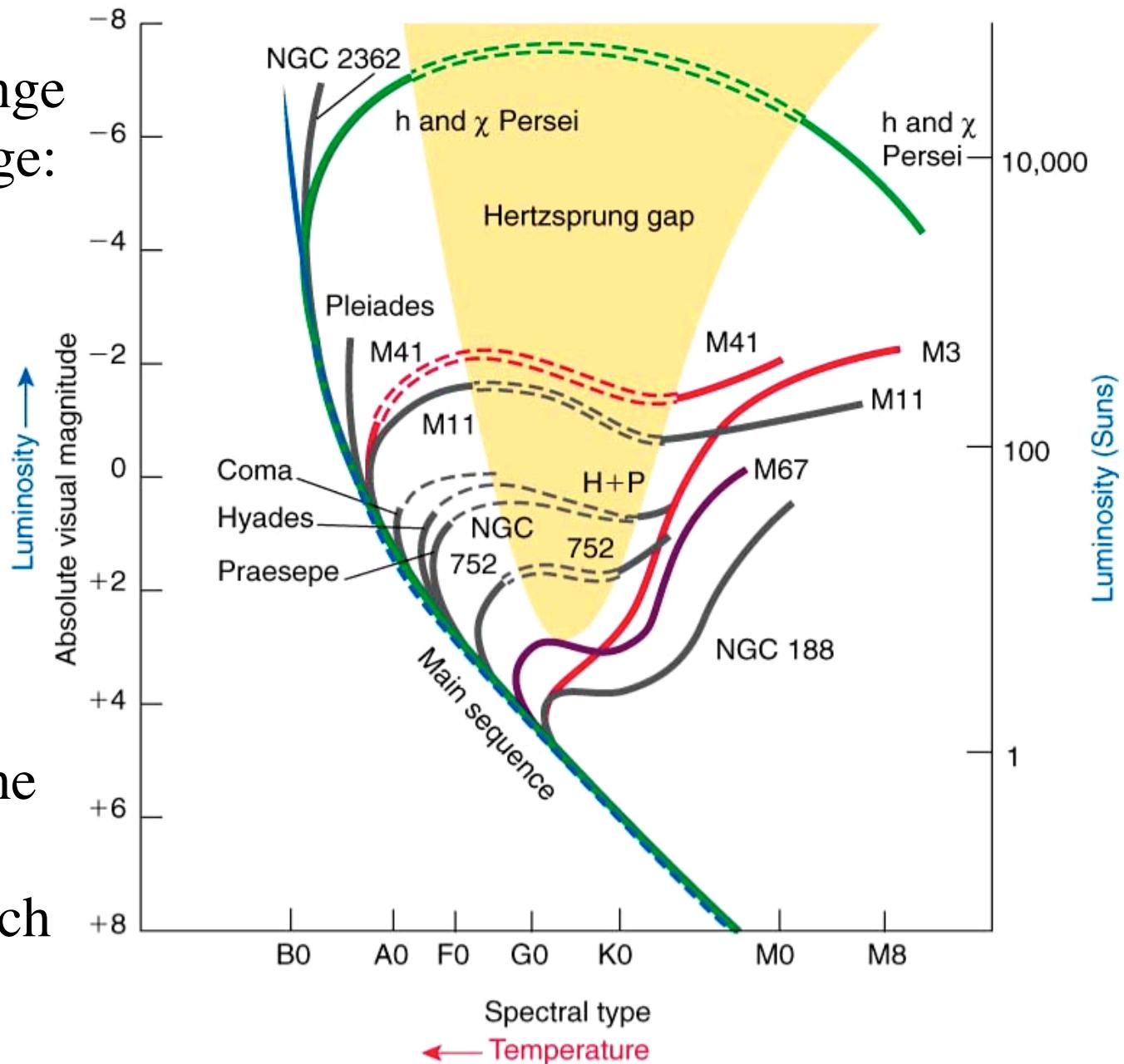
Testing Stellar Evolution

- The problem: stellar evolution happens on billion-year time scales
- The solution: use HR Diagrams of *star clusters* with a wide range of ages
 - Clusters contain 100's to 1000's of stars with a broad mass range
 - All stars are at the same distance, so it is easy to measure their relative luminosities
 - They have the same age, have the same chemical composition
- *Each cluster thus provides a snapshot* of what stars of different masses look like at the same age and composition (coeval populations)

Open Clusters: HR Diagrams

A systematic change
with the cluster age:

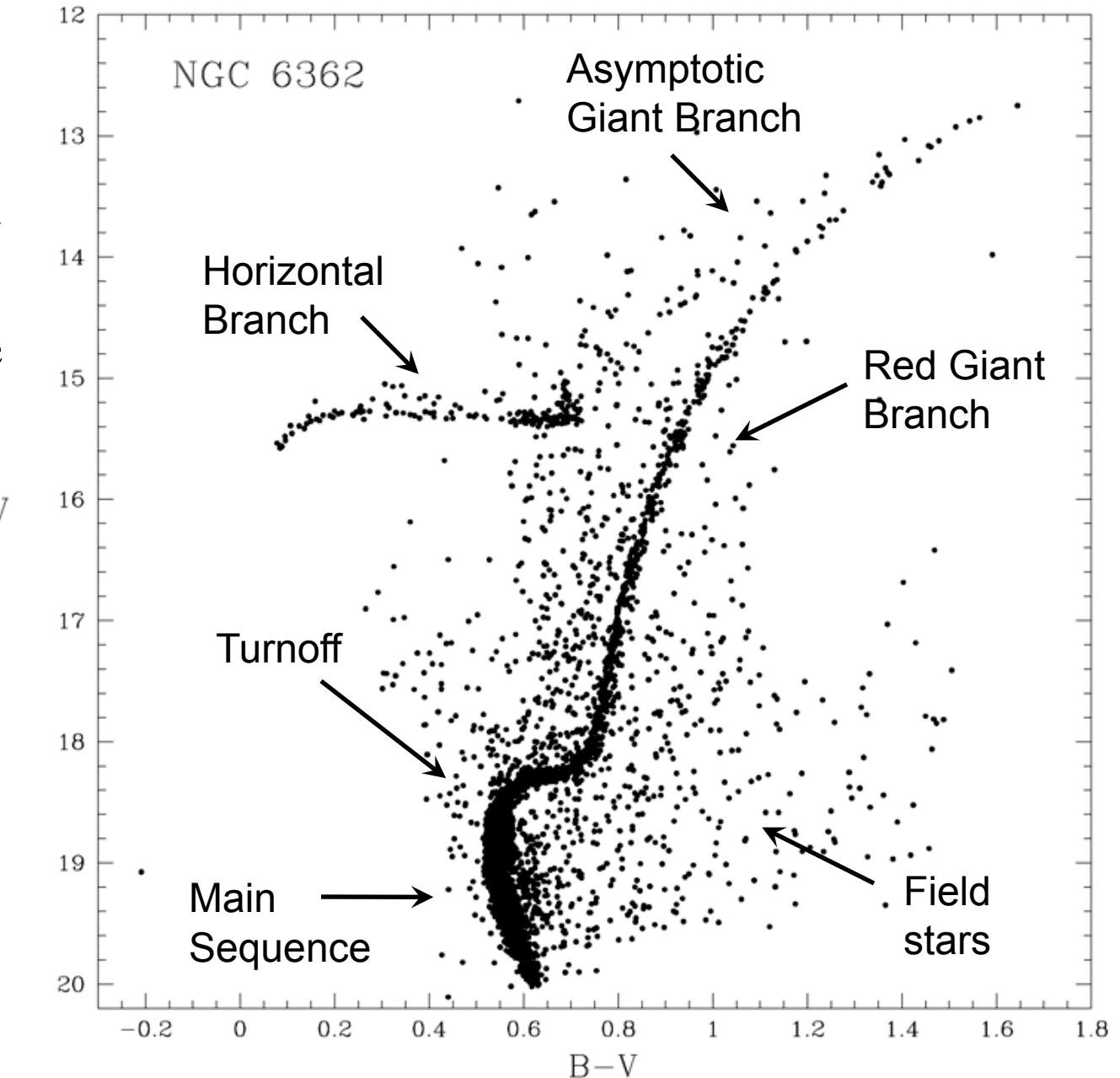
The MS turnoff
moves to lower
luminosities and
temperatures as the
cluster ages, and
the red giant branch
develops



Globular Clusters: HR Diagrams

Stars above the turnoff have evolved away, and other branches represent the more advanced stages of stellar evolution

We can use that to estimate the ages of the globular clusters, and thus of the Galaxy

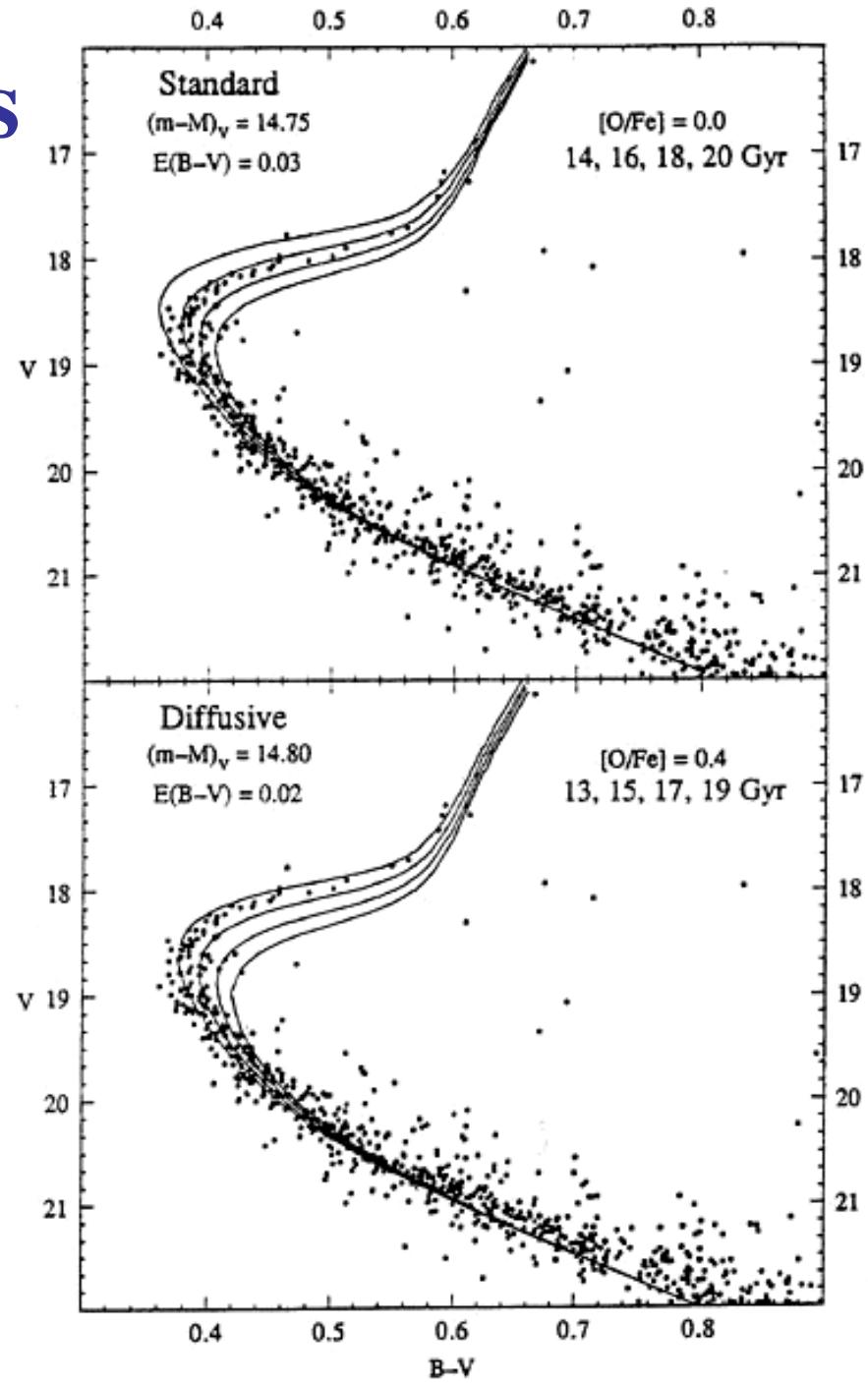


Globular Cluster Ages

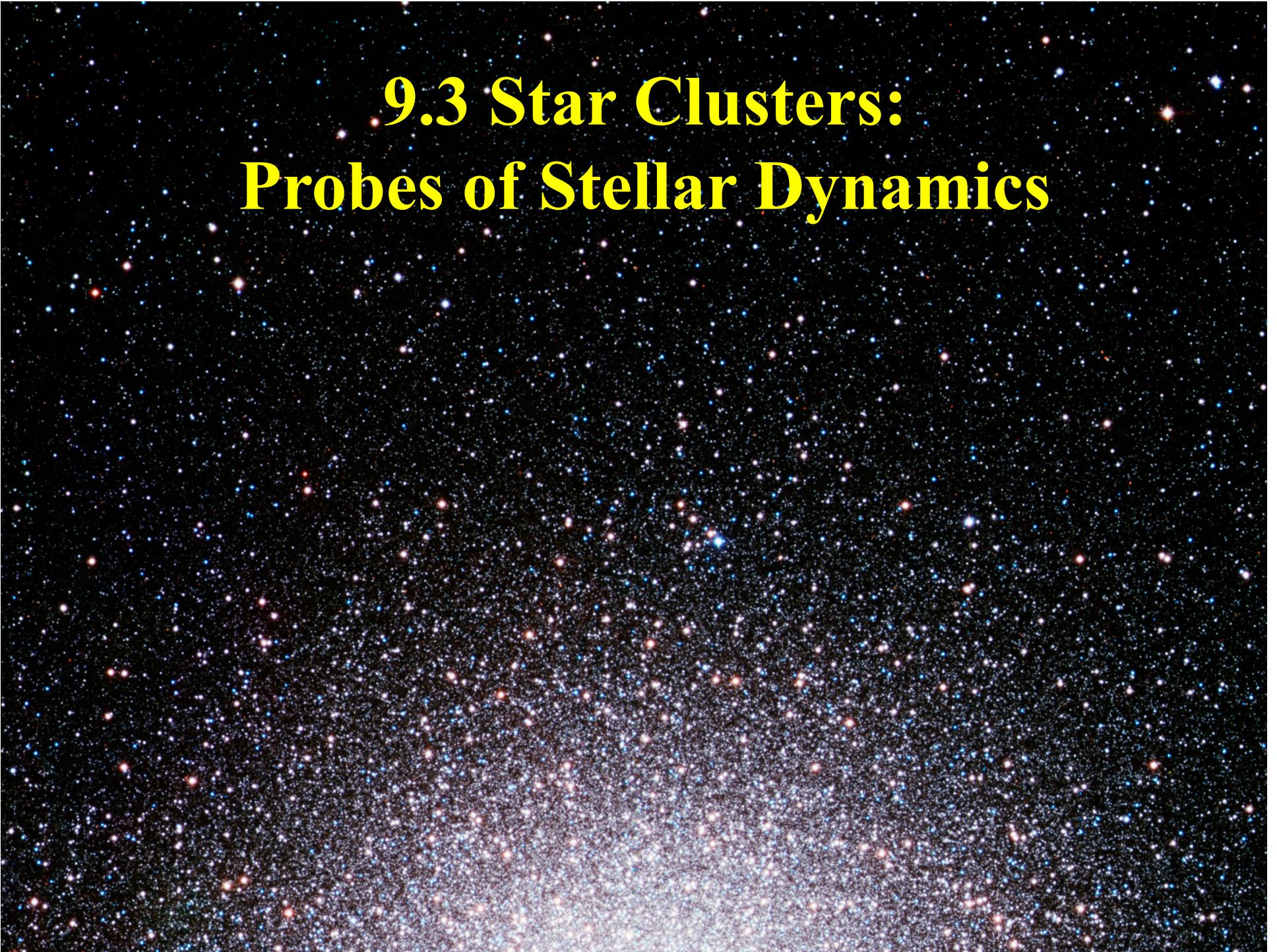
Lines: *Iochrones* =
theoretical models of the main
sequence and the red giant
branch (and others), at a given
age, for a given chemical
composition

Modern value for the mean
globular cluster age in our
Galaxy:

$$\text{Age} = 12.3^{+2.1}_{-2.5} \text{ Gyr}$$



9.3 Star Clusters: Probes of Stellar Dynamics



Dynamical Evolution of Star Clusters

1. Internal processes:

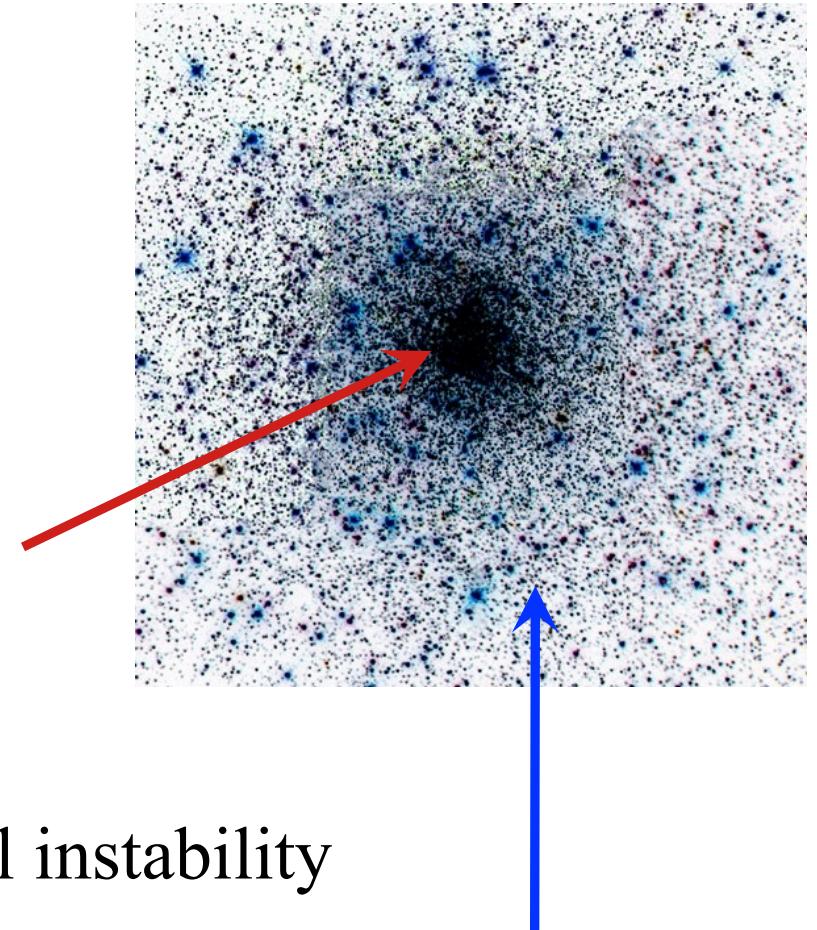
- *Dynamical relaxation*: stars exchange energies in 2-body interactions, over a relaxation time

Stars in the core have lower energies and sink to the bottom of the potential well

- Core collapse, or gravothermal instability

2. External processes:

- Tidal shocks and evaporation



Stars with higher energies can reach large radii

Dynamical Time Scales

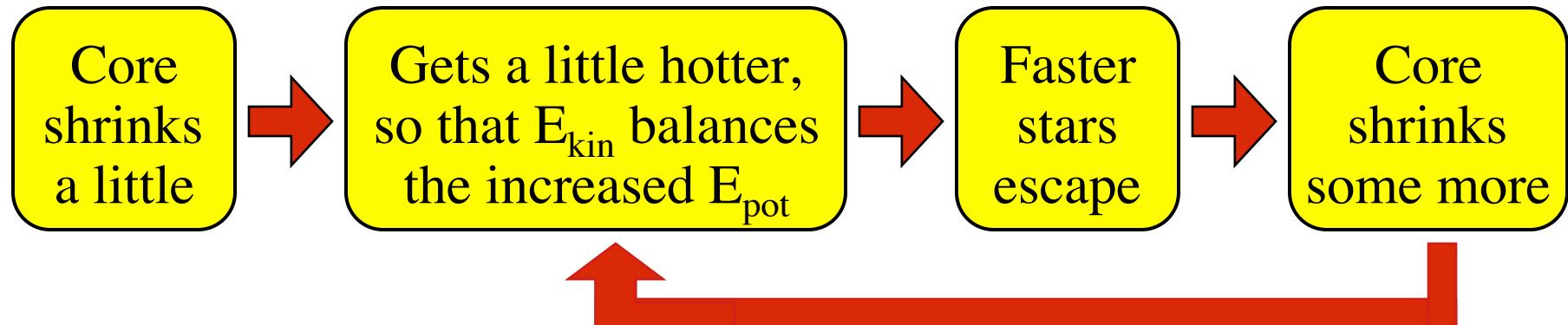
	N	R (pc)	V_{total} (km/s)	$t_{cross} = R/V$ ($\times 10^6$ yr)	t_{relax} (yr)
Open cluster	100	2	0.5	4	8×10^6
Globular cluster	10^5	4	10	0.4	4×10^8
E galaxy	10^{12}	10	600	20	1×10^{17}

Open clusters: $t_c \sim t_r < t_e \rightarrow$ quickly dissolved

Globular clusters: $t_c \ll t_r \ll t_e \rightarrow$ a variety of dynamical evolution states must be present

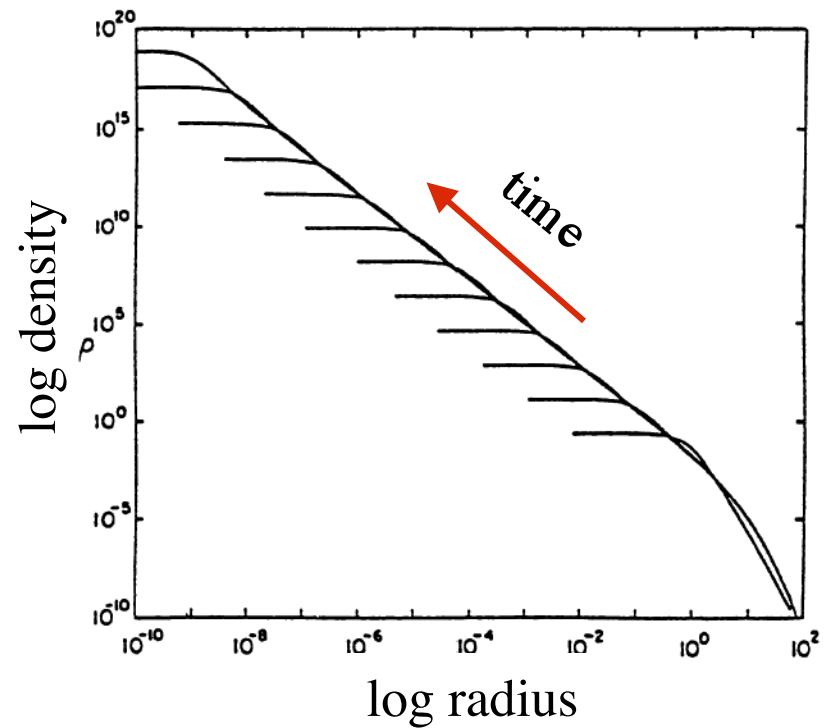
Elliptical galaxies: $t_c \ll t_r \sim t_e \rightarrow$ dynamical evolution not driven by 2-body relaxation

Core Collapse, aka The Gravothermal Catastrophe



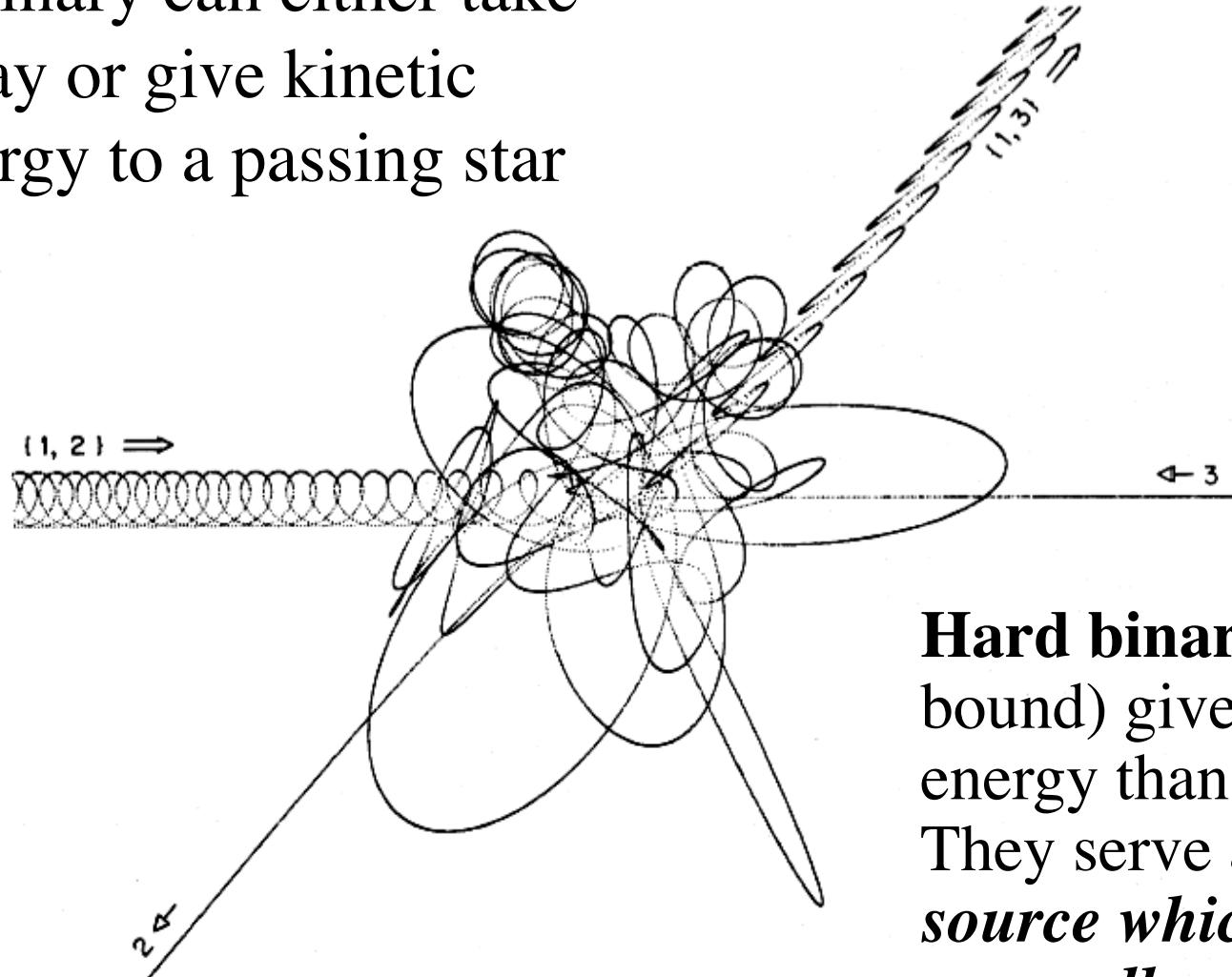
The only way to arrest the collapse is to provide *a source of energy* in the center, to replace the escaped heat. In the case of (proto)stars, this is accomplished by thermonuclear reactions.

In the case of globular clusters, it is accomplished by ***hard binaries***.



Binaries as a Source of Energy

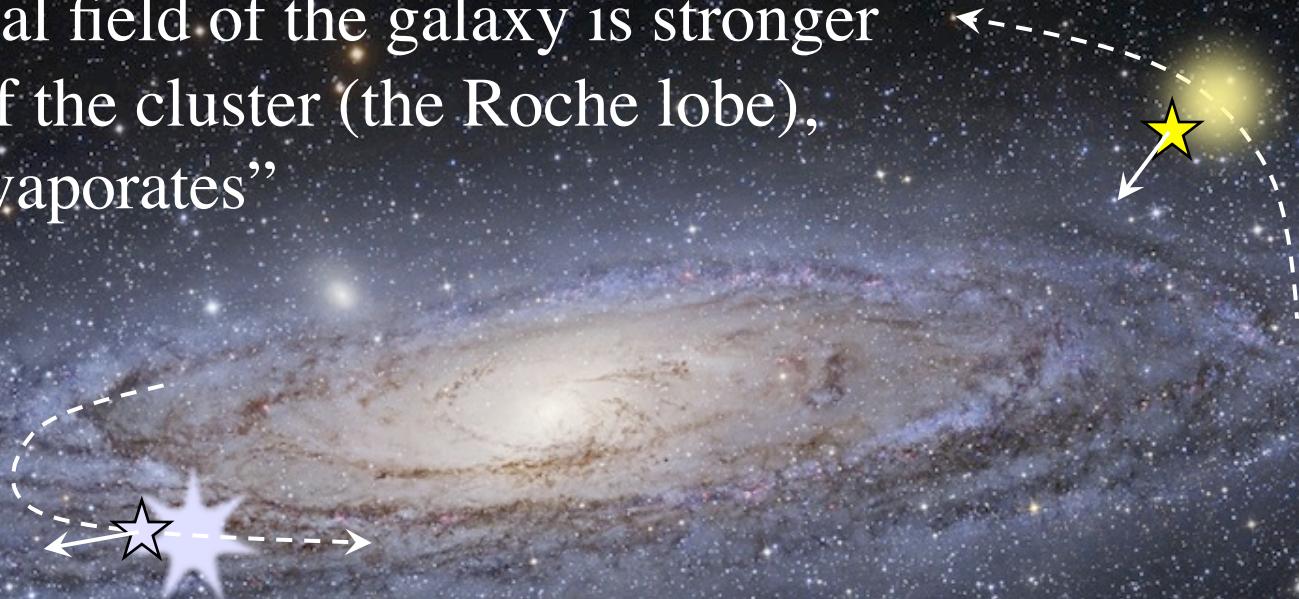
A binary can either take away or give kinetic energy to a passing star



(Numerical simulation by
P. Hut and J. Bahcall)

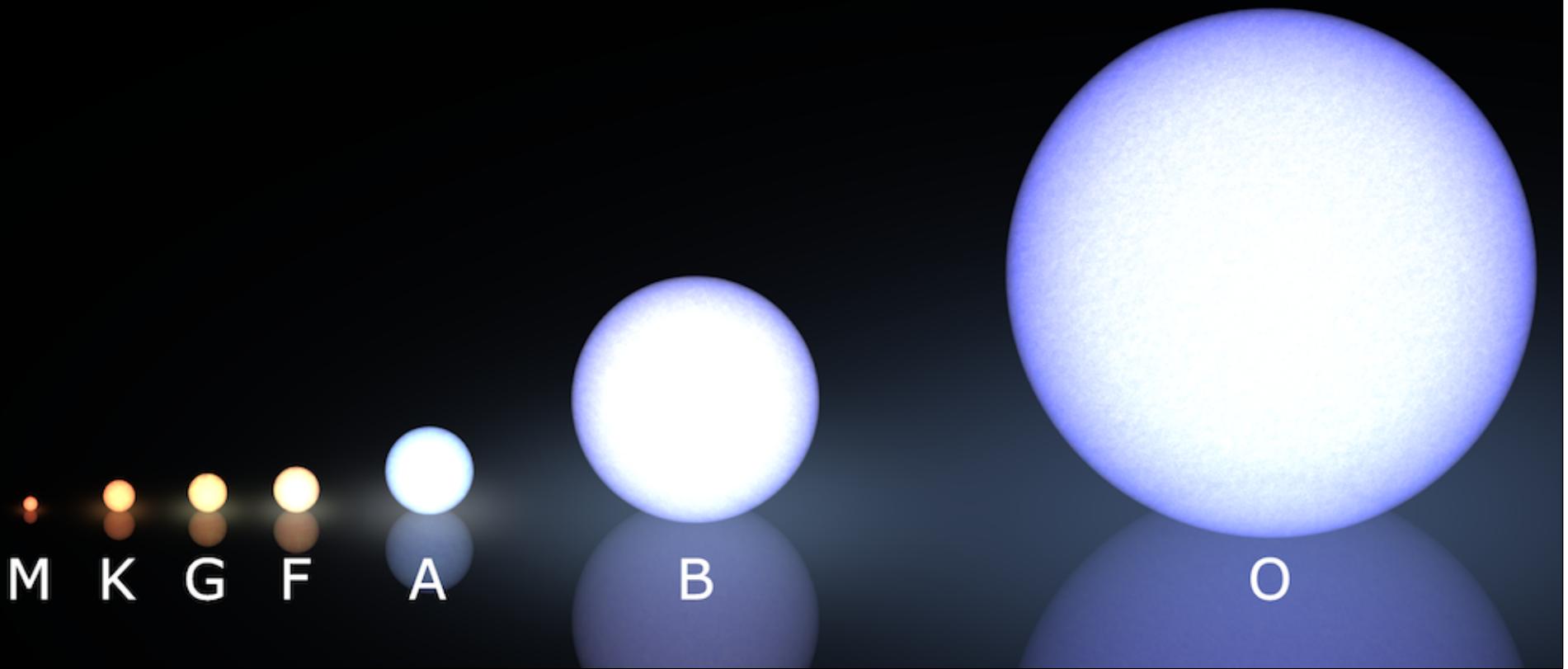
Hard binaries (tightly bound) give away more energy than they absorb. They serve as the *energy source which arrests the core collapse* and stabilize the cluster

Globular clusters move in the tidal field of the galaxy.
When a star crosses the boundary where the gravitational field of the galaxy is stronger than that of the cluster (the Roche lobe), the star “evaporates”



The same thing happens to stars in open clusters as they pass by the giant molecular clouds in the disk

9.4 The Main Sequence



Main Sequence (MS) and the Range of Stellar Masses

- MS is defined as the locus where stars burn H into He in their cores
- That is where they spend most of their lifetime
- It is *a sequence of stellar masses* – by far the most dominant parameter that determines stellar properties
- The lower mass end is set by the objects which cannot reach the necessary [T, ρ] to ignite fusion, because of their ***low mass*** ($M_\star < 0.08 M_\odot$): **brown dwarfs** (new spectral types: L, T, Y)
- The ***high-mass end*** of the stellar family is set by the **Eddington limit**

Evolution on the Main Sequence

Star burns H in core, core composition slowly changes from H to He. Small changes in the external properties (L, T_e, R)

Main-sequence lifetime is strongly mass-dependent, since the more massive stars:

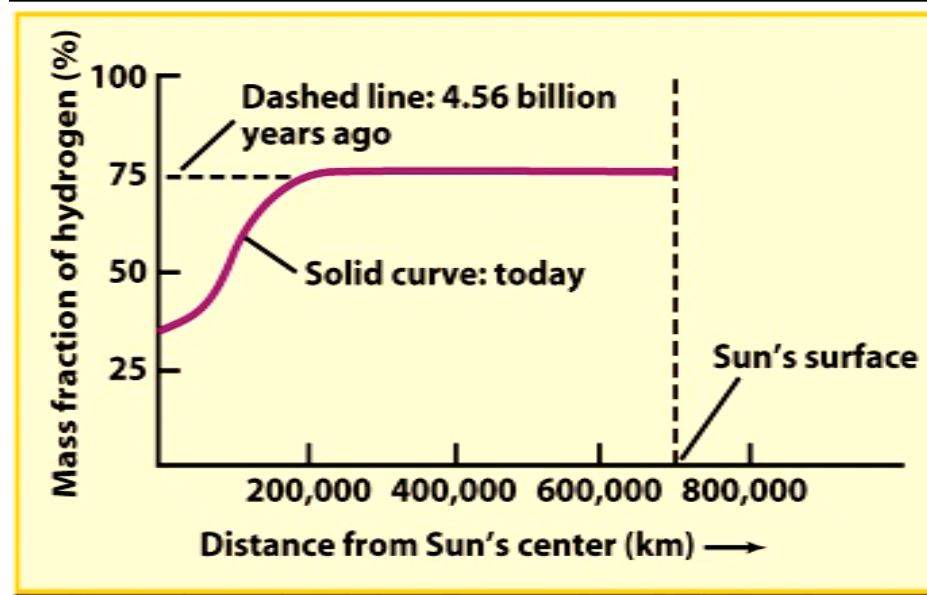
- sustain higher core temperatures
- have higher rates of nuclear fusion
- are more luminous and exhaust H fuel more quickly

$$L \propto M^{3.5} \rightarrow t_{ms} \propto M^{-2.5}$$

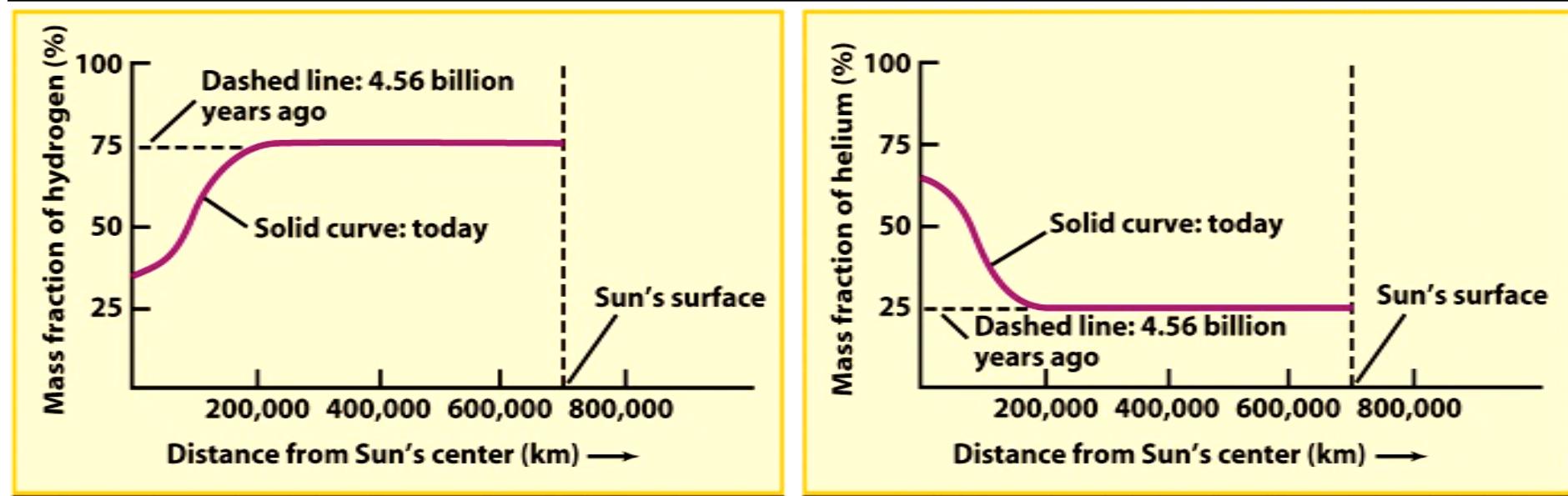
Star leaves the main sequence when it stops burning hydrogen in the core that is now pure He (but it continues burning it in a shell around the core). This leads to *expansion* of the envelope, and the formation of a **red giant**

Stellar Lifetimes on the Main Sequence

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



(a) Hydrogen in the Sun's interior

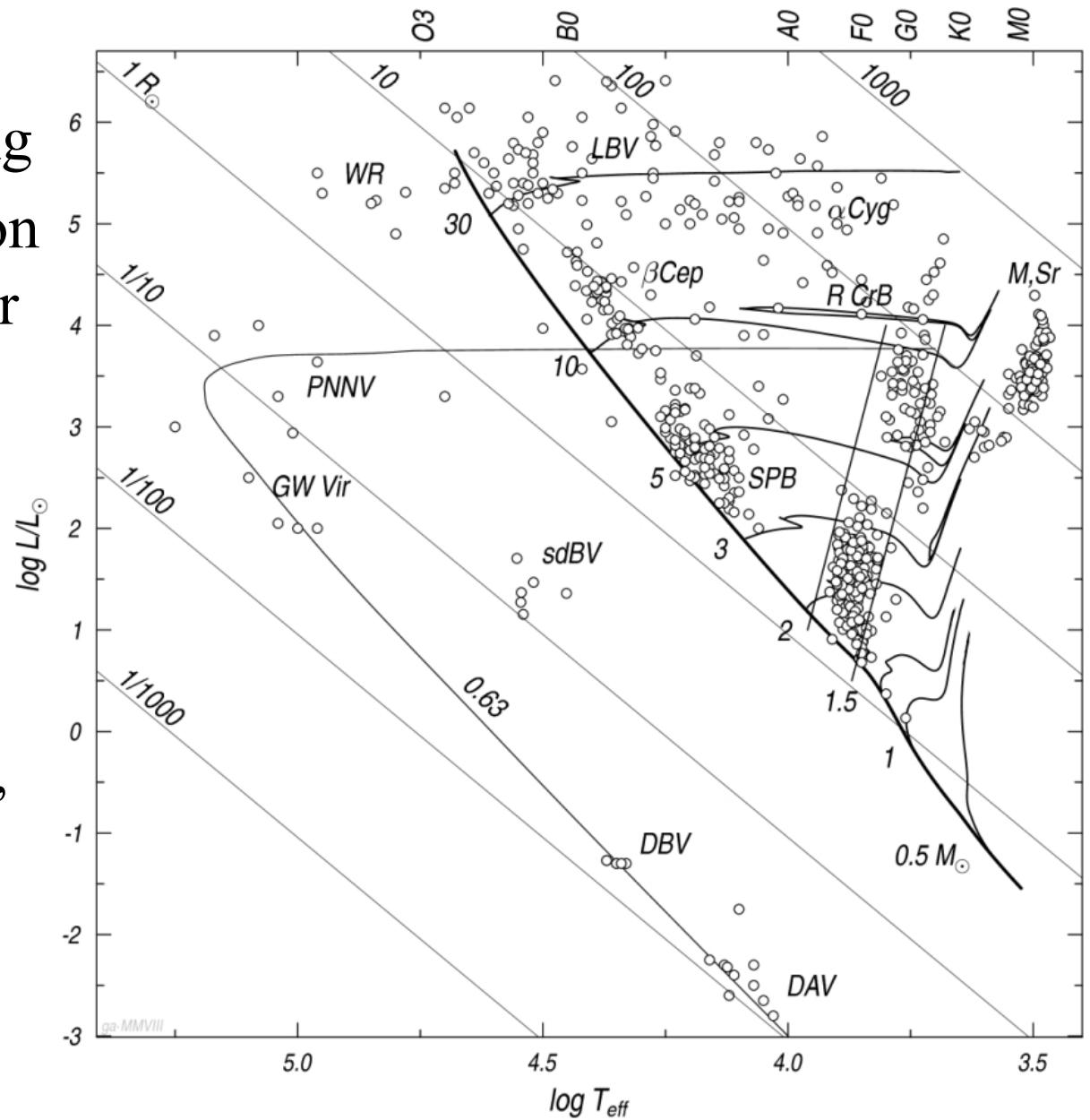


(b) Helium in the Sun's interior

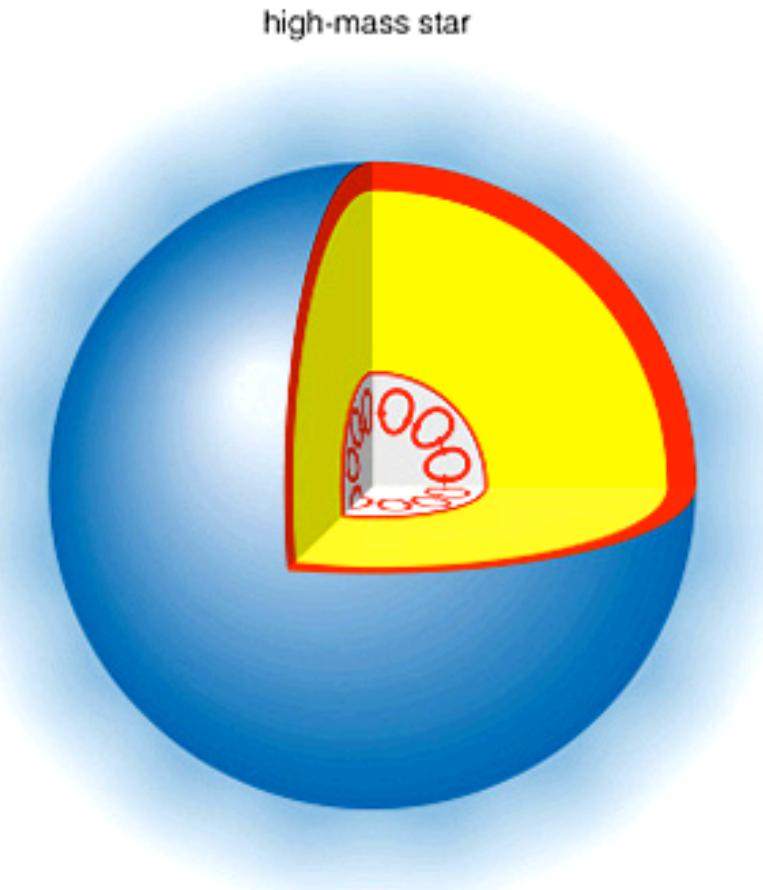
Global Stellar Properties

Mass essentially determines everything
Chemical composition has a relatively minor effect

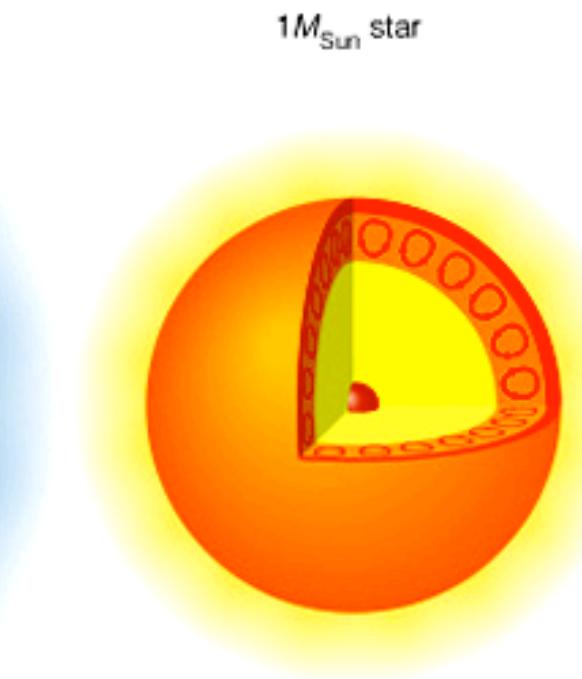
Position of a star in the (L,T) plane also determines its radius, through the Stefan-Boltzmann formula, and thus also the density



Differences in Stellar Structures Regarding the Energy Transport



Convective core,
radiative envelope



Radiative core,
convective envelope



Fully
convective

Eddington Limit

Electrons/ions at a stellar surface feel radiation pressure that is proportional to luminosity, and that can drive a *stellar wind*

When the radiation pressure matches the gravitational pull of the star, we have a limiting, **Eddington luminosity:**

$$L = \frac{4\pi G c m_p}{\sigma_e} M$$

A more luminous star at a given mass would blow itself apart.

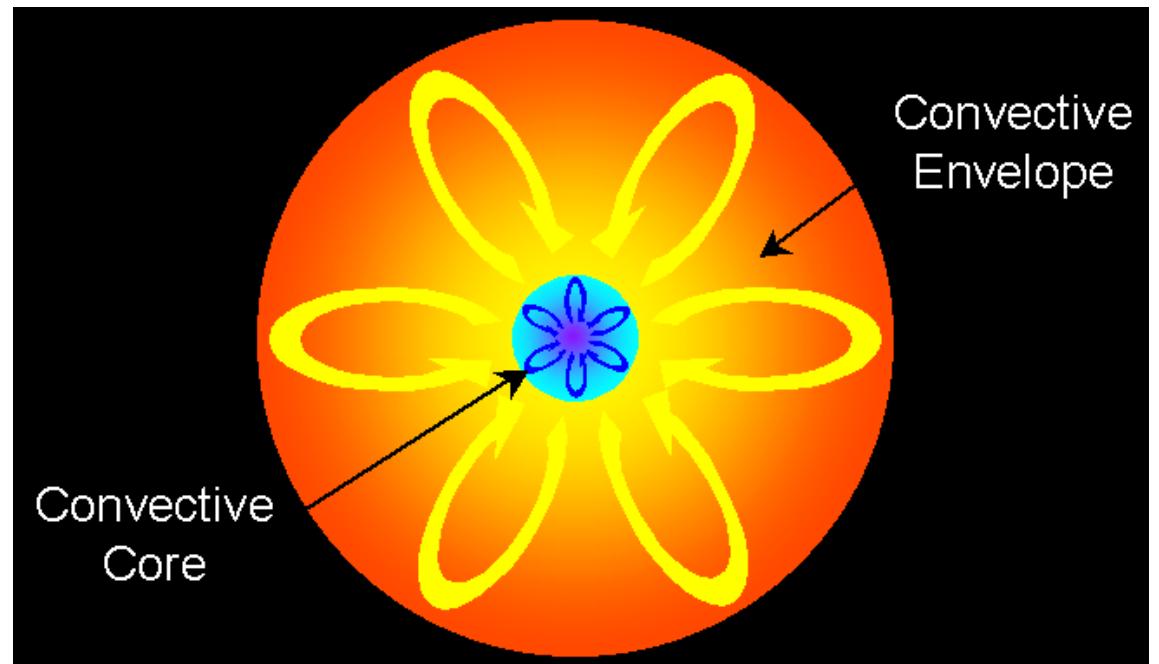
$$= 1.26 \times 10^{38} \left(\frac{M}{M_{sun}} \right) \text{erg s}^{-1}$$

This is the *maximum luminosity* which an isotropically emitting source with a mass M could have

Invert the formula: $M_E = 8 \times 10^5 \left(\frac{L}{10^{44} \text{ erg s}^{-1}} \right) M_{sun}$

Very Low Mass Stars: Red Dwarfs

- Mass $< 0.4 M_{\odot}$
- Their structure is all convection zone, H and He is mixed throughout the star
- The star burns H slowly
- Will never build up a He core, and never ignite He
- Could perhaps survive on the MS for a 100 Gyr!
- Then just fade as a White dwarf



The End of the MS Phase

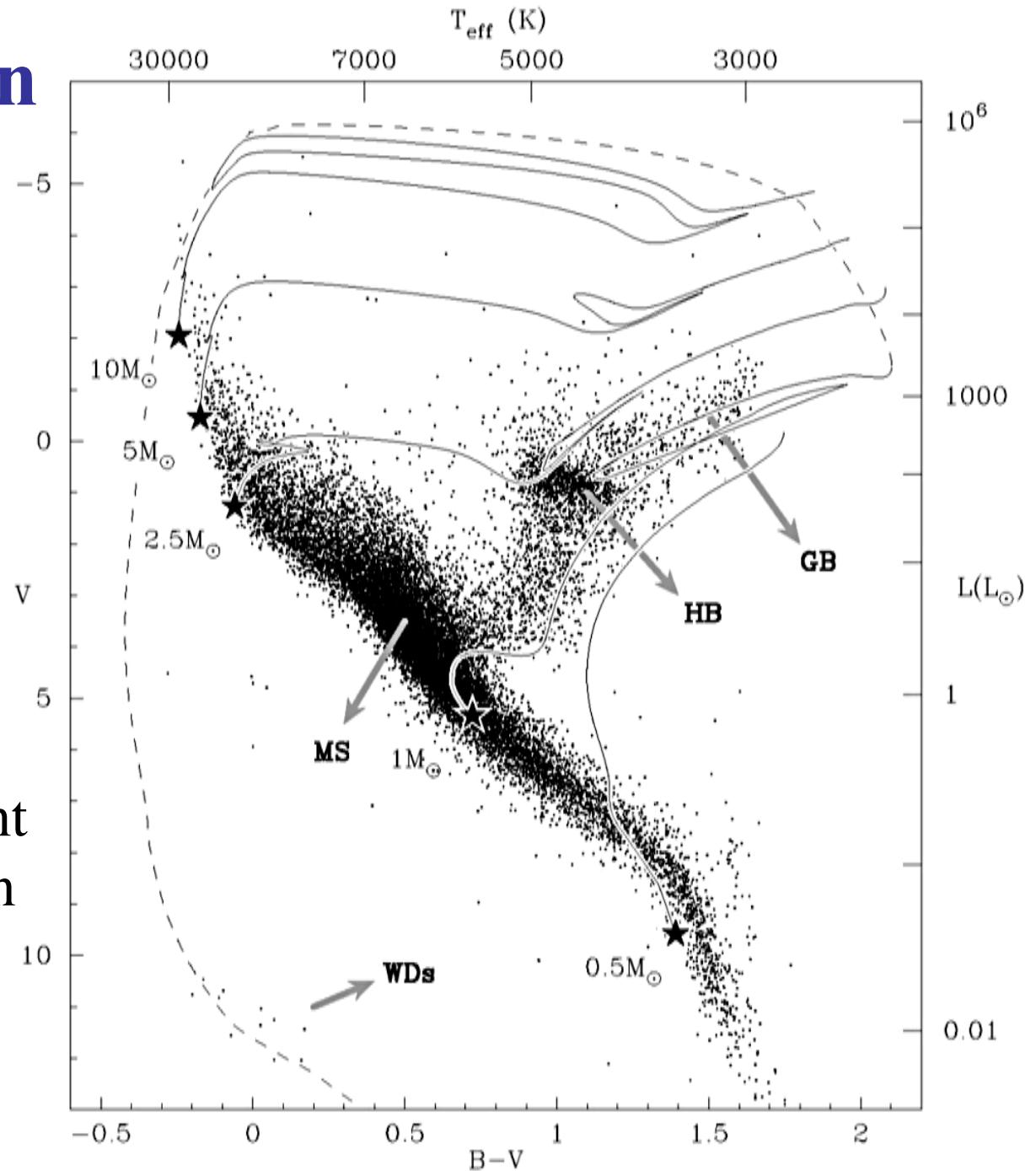
- On the MS, a star is in a hydrostatic equilibrium, and its core is sufficiently hot to fuse H into He
- Now the star has two chemically distinct zones, a core of inert He surrounded by an H envelope - the core of a MS star is not sufficiently hot for He burning
- When the core becomes pure He, a new evolutionary phase starts - the ascent to the ***Red Giant Branch*** (RGB)
- Without energy generation, the core cannot support itself against gravitational collapse and so it begins to shrink; as it collapses it heats up
- This heat is transferred to a thin shell of H around the core which reaches a temperature in which H fusion can occur

9.5 Stellar Evolution After the Main Sequence

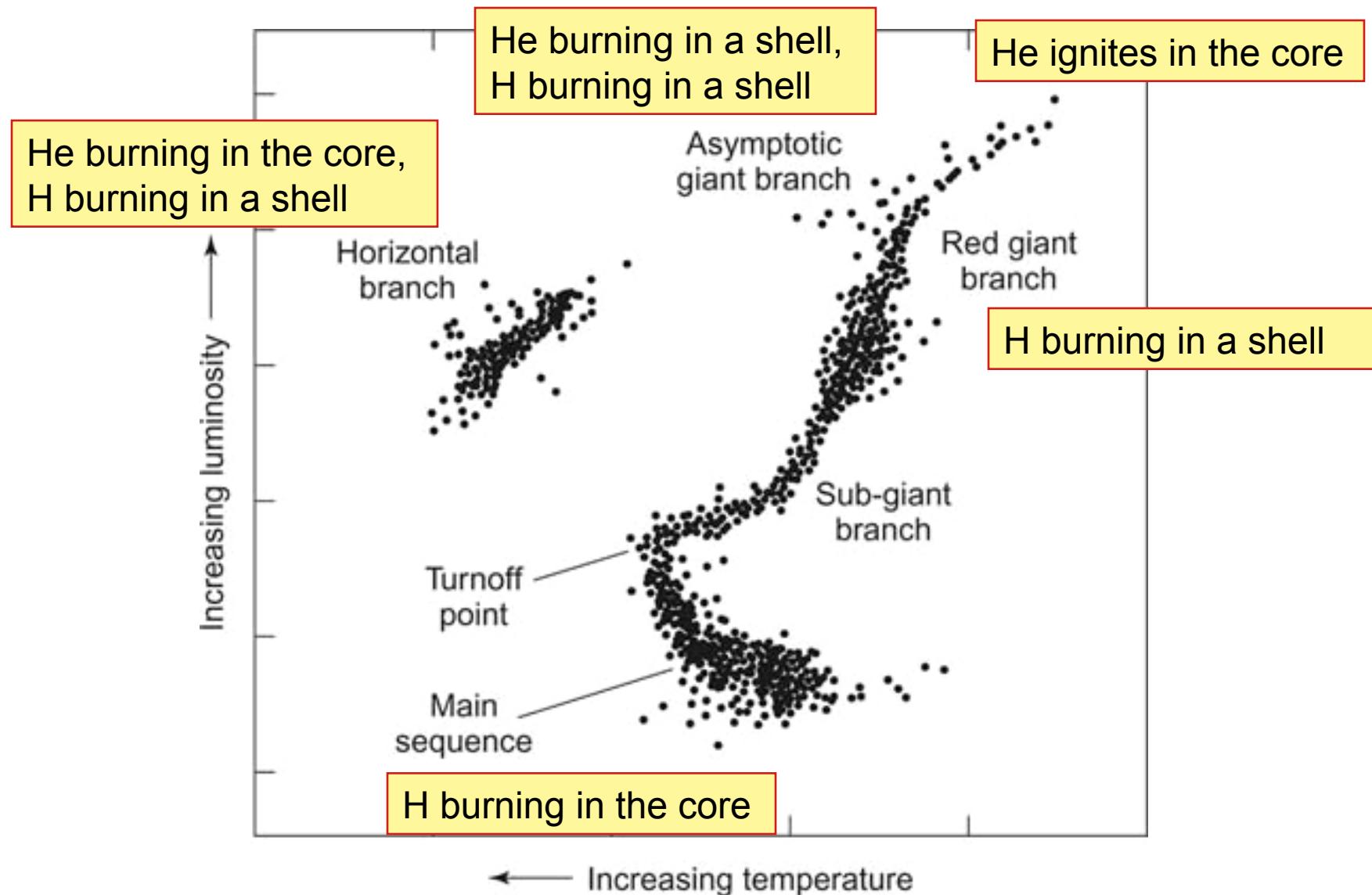


Stellar Evolution Off the Main Sequence

Stars of different masses follow different *evolutionary tracks* on the HR diagram



Stellar Evolution is a Sequence of Different Thermonuclear Reactions

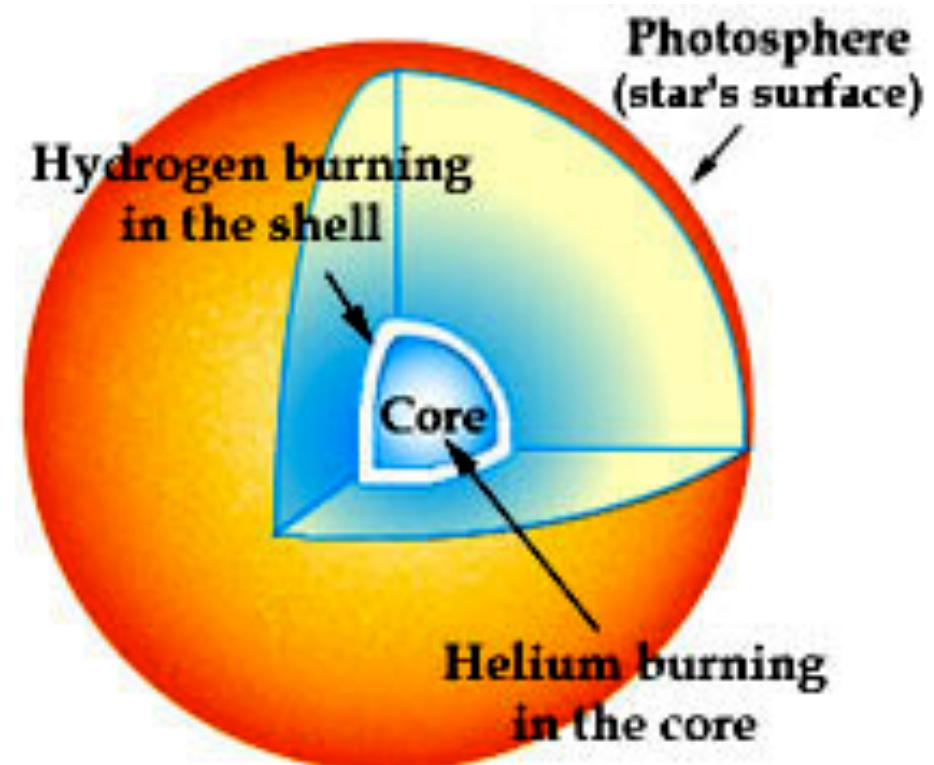


Becoming a Red Giant

- As the core continues to collapse, the temperature in the H fusing shell continues to rise and thus the luminosity in the shell increases as does the pressure
- The entire star is no longer in a hydrostatic equilibrium, and the envelope begins to expand
- As they expand these outer layers cool - the star becomes redder, while its luminosity increases: the star slowly ascends the RGB
- Our Sun will swell to about the size of the Earth's orbit
- This imbalance will continue until the star again finds a source of core energy generation, i.e., He fusion

Structure of a Red Giant

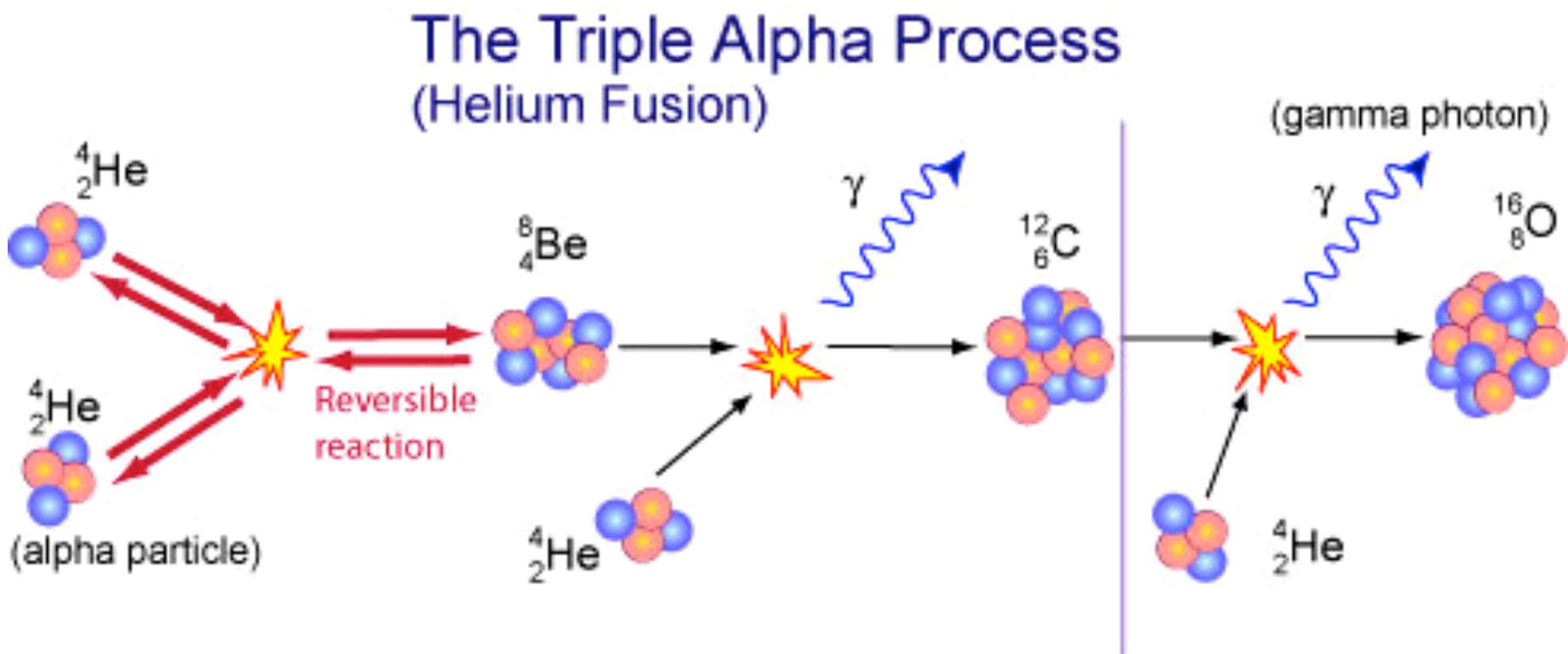
- Now the core has heated to $T \sim 10^8$ K, which is the threshold temperature for the fusion of He into C
- The star is in a quasi-static equilibrium. The lifetime of a star as a Red Giant is about 10% of its MS lifetime
- The luminosity generated by the core fusion of He into C is far greater than the shell luminosity associated with the fusion of H into He



Helium Flash: The End of the RGB

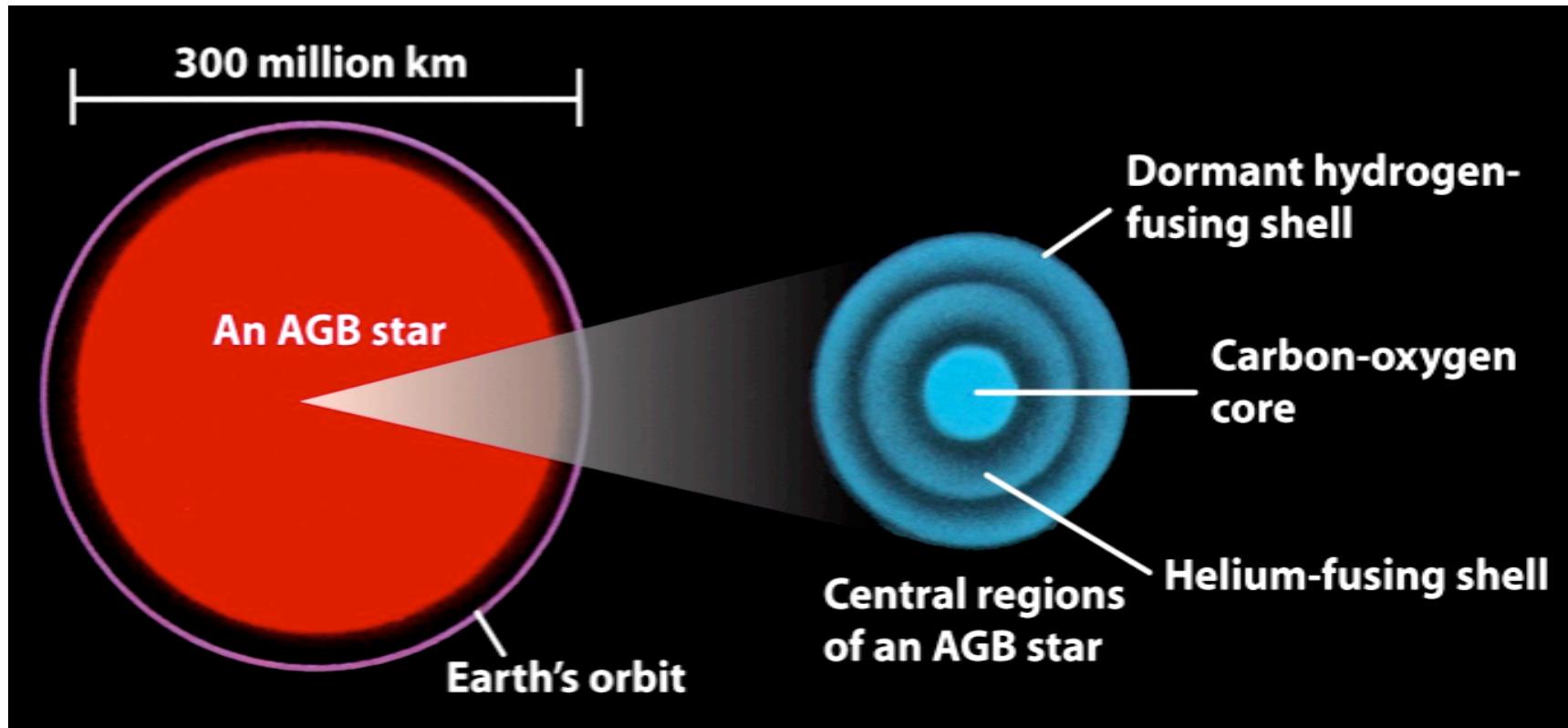
- H fusion leaves behind He ash in the core of the star which cannot begin to fuse until the temperature of the core reaches 100 million K. How a star begins He fusion depends on its mass:
- $M > 3 M_{\odot}$ stars contract rapidly, their cores heat up, and He fusion begins gradually
- Less massive stars evolve more slowly and their cores contract so much that degeneracy occurs in the core
- When the temperature is hot enough He fusion begins to make energy and the T rises, but pressure does not increase due to degeneracy
- Higher T increases He fusion even further resulting in a runaway explosion: the **Helium Flash** which for a few minutes can generate more luminosity than an entire galaxy. The flash does not destroy the star: the envelope absorbs the energy

The Next Step: Burning Helium Into Carbon



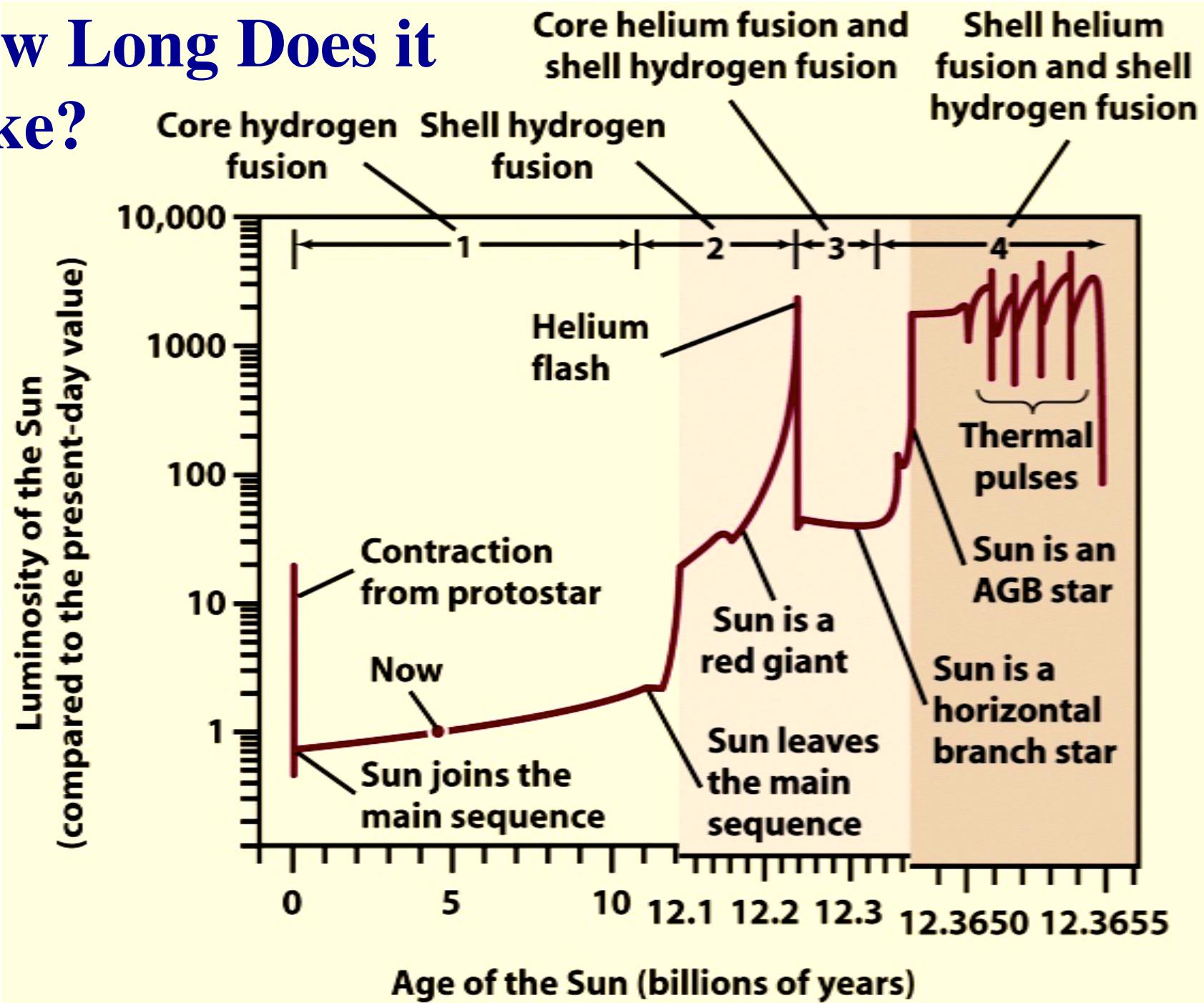
- Requires much higher temperatures, $T \sim 10^8 \text{ K}$
- Enabled by the “exact right” energy resonance for carbon

Dredge-ups bring the products of nuclear fusion to a giant star's surface



- As a low-mass star ages, convection occurs over a larger portion of its volume
- This takes heavy elements formed in the star's interior and distributes them throughout the star

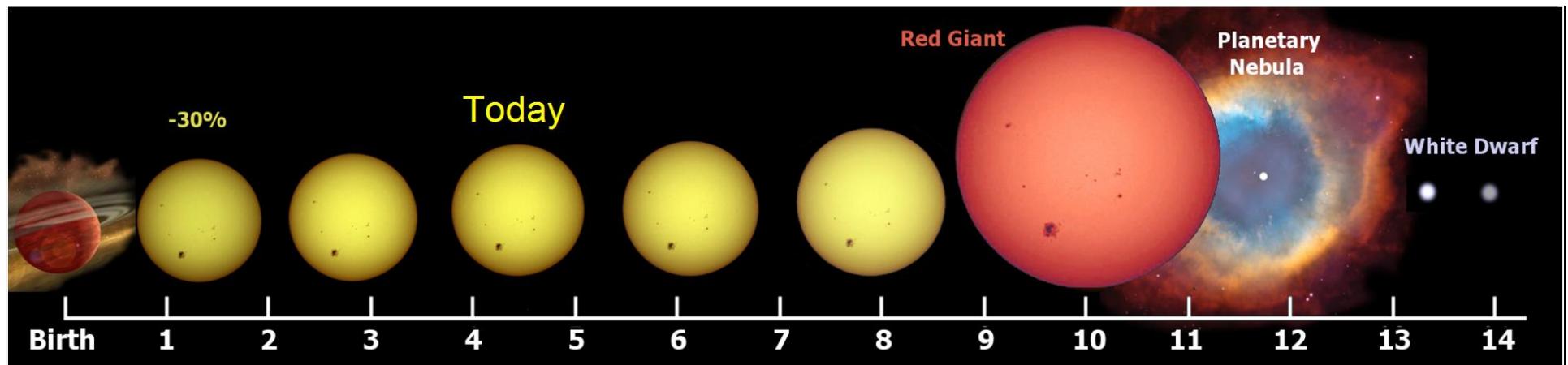
How Long Does it Take?



The End Phases of Stellar Evolution

- The evolution and eventual fate of stars critically dependent on their mass:
- Stars with initial masses of *less* than $\sim 8 M_{\odot}$ end as ***white dwarfs***. The star sheds its RG envelope, which becomes a ***planetary nebula***, and the inert, degenerate core cools passively
- Stars with initial masses *greater* than $\sim 8 M_{\odot}$ explode as ***supernovae***. The stellar remnants are ***neutron stars*** or ***black holes***

The fate of our Sun

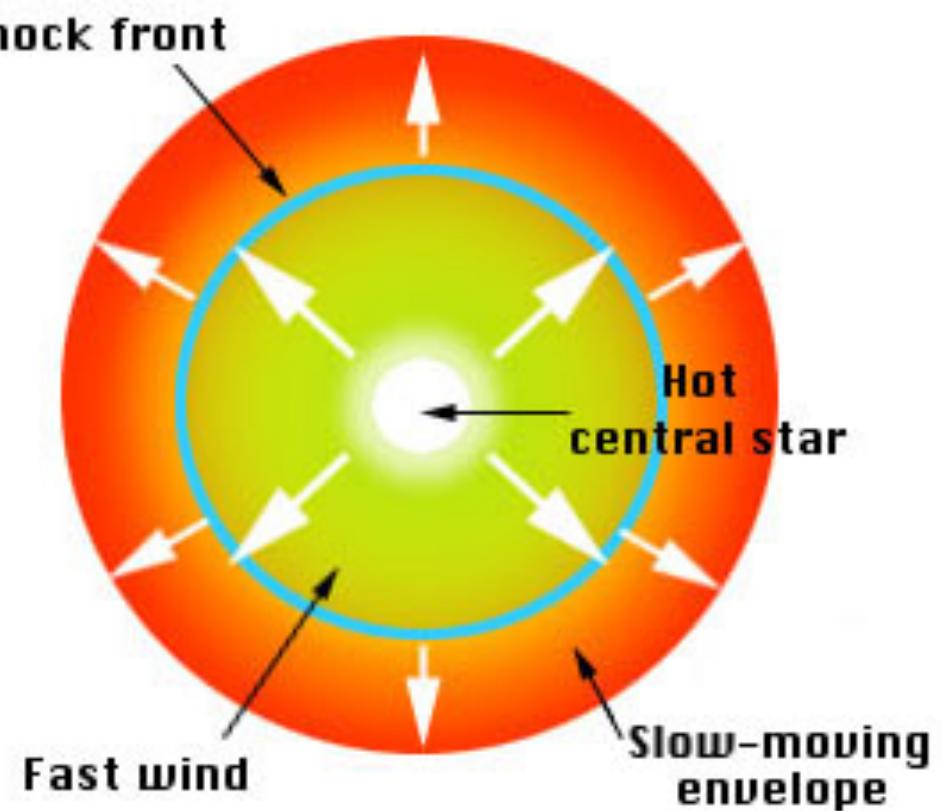


Planetary Nebula Formation

- A RG brightens by a factor of between 1,000 and 10,000. The outer, hydrogen-rich envelope swells up to a few au radius, with $T \sim 2,000 - 3,000$ K
- A strong stellar wind begins to blow from the star's surface (akin to the Sun's solar wind, but much stronger), and, in the course of the star's RG life, carries away most of the H envelope
- During the final shedding of its envelope, when the mass loss is the greatest, the star becomes unstable and pulsates, with periods \sim few months to > 1 yr. Such stars are called ***long-period variables***.
- The envelope material ejected by the star forms an expanding shell of gas that is known as a ***planetary nebula (PN)***

Formation of Planetary Nebula

Planetary nebulae typically have masses $\sim 0.2 M_{\odot}$, although some are considerably more massive. They expand $V \sim 10 - 20$ km/s, and plow into the surrounding *interstellar medium*, contributing to its chemical enrichment.

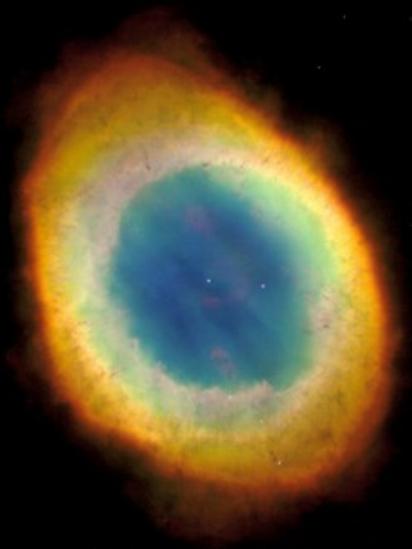


Their central stars (degenerate cores of progenitors) provide photoionization for the nebulae.

Planetary Nebulae Imaged by the HST



Eskimo Nebula



Ring Nebula



Necklace Nebula



Spirograph Nebula (IC 418)

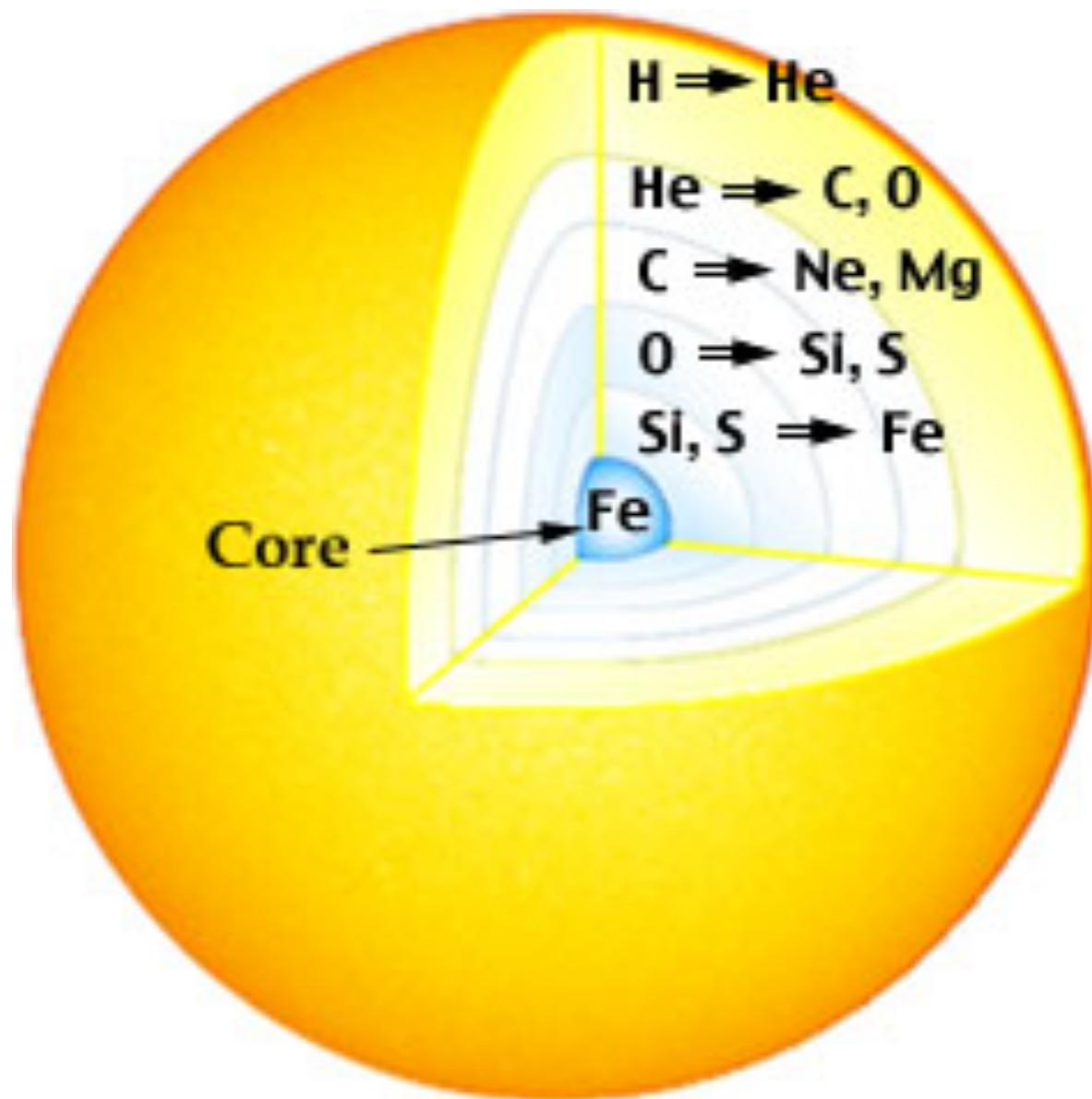


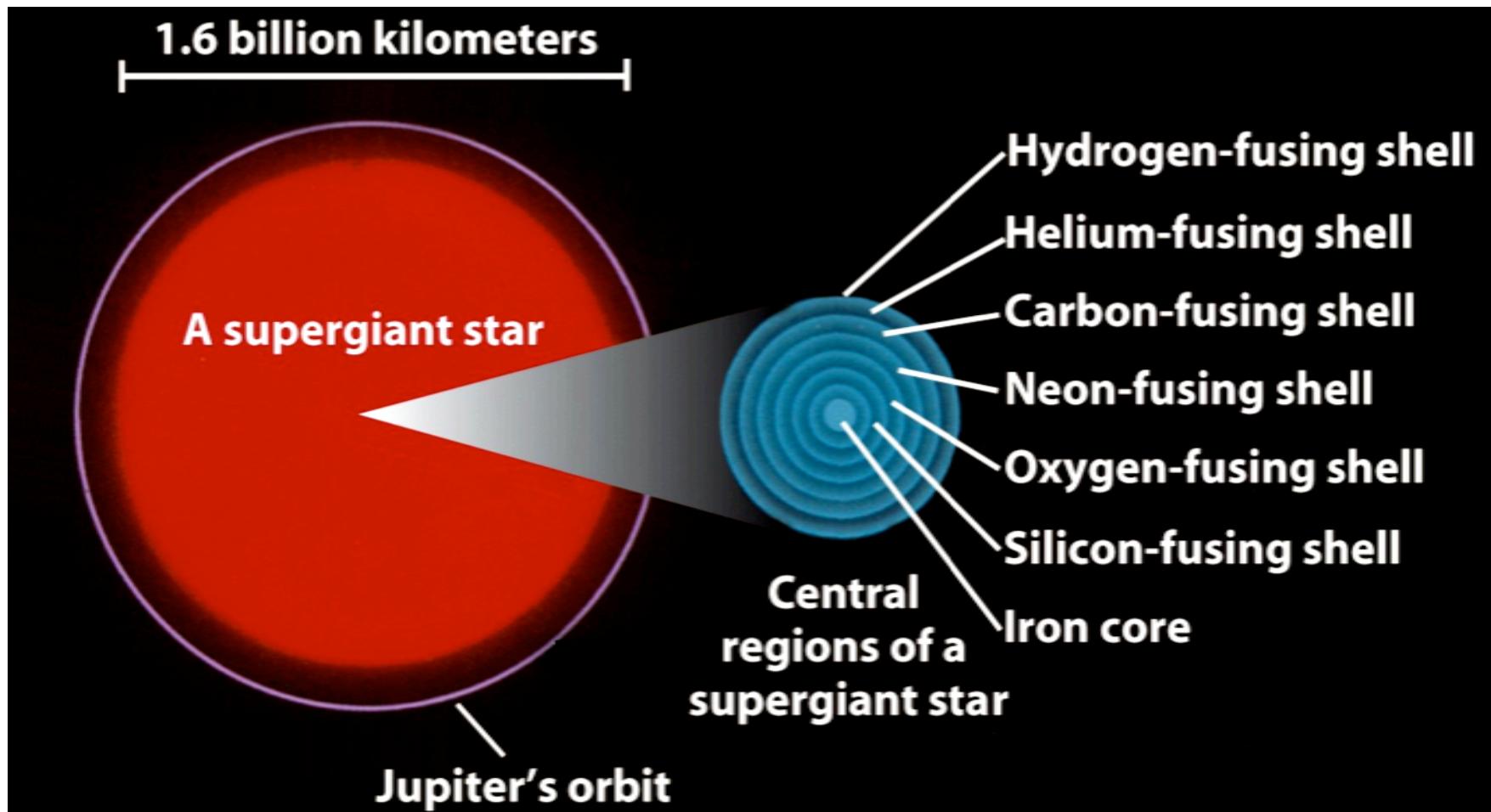
Cat's Eye Nebula



Hour Glass Nebula

Highly Evolved Star Structure





- In the last stages, a high-mass star has an Fe-rich core surrounded by concentric shells hosting the various thermonuclear reactions
- The sequence of thermonuclear reactions stops here, because the formation of elements heavier than Fe requires an input of energy rather than causing energy to be released

Rapid Final Stages of a Massive Star Evolution

- A high mass star undergoes an extended sequence of thermonuclear reactions in its core and shells: C fusion, Ne fusion, O fusion, and Si fusion, all the way to Fe

Stage	Core temperature (K)	Core density (kg/m ³)	Duration of stage
Hydrogen fusion	4×10^7	5×10^3	7×10^6 years
Helium fusion	2×10^8	7×10^5	7×10^5 years
Carbon fusion	6×10^8	2×10^8	600 years
Neon fusion	1.2×10^9	4×10^9	1 year
Oxygen fusion	1.5×10^9	10^{10}	6 months
Silicon fusion	2.7×10^9	3×10^{10}	1 day
Core collapse	5.4×10^9	3×10^{12}	$\frac{1}{4}$ second
Core bounce	2.3×10^{10}	4×10^{15}	milliseconds
Explosive (supernova)	about 10^9	varies	10 seconds

From Birth to Death

