From Massive Cores to Massive Stars

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Star Formation, Then and Now
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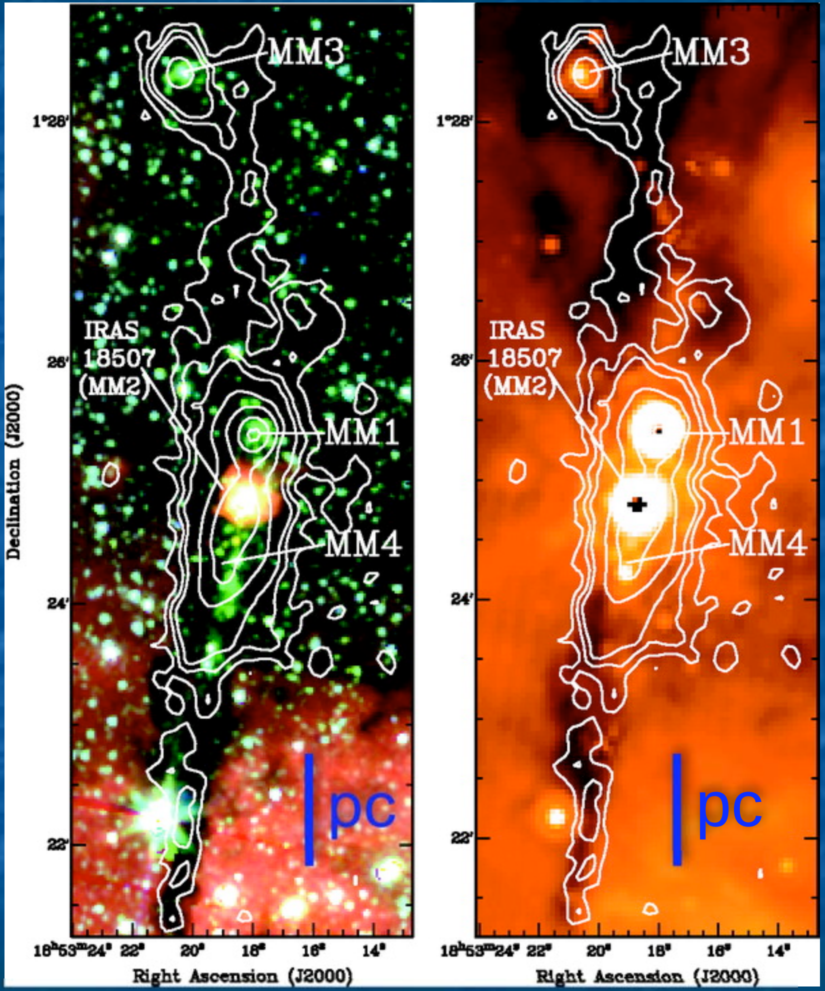
Talk Outline

- Massive cores
- From core to star
  - Fragmentation
  - Disk Formation and Accretion
  - Radiation Pressure Feedback (deferred to Richard Klein’s talk)
  - Competitive Accretion
- Final summary

Sites of Massive Star Formation

(Plume et al. 1997; Shirley et al. 2003; Rathbone et al. 2005; Yonekura et al. 2005)

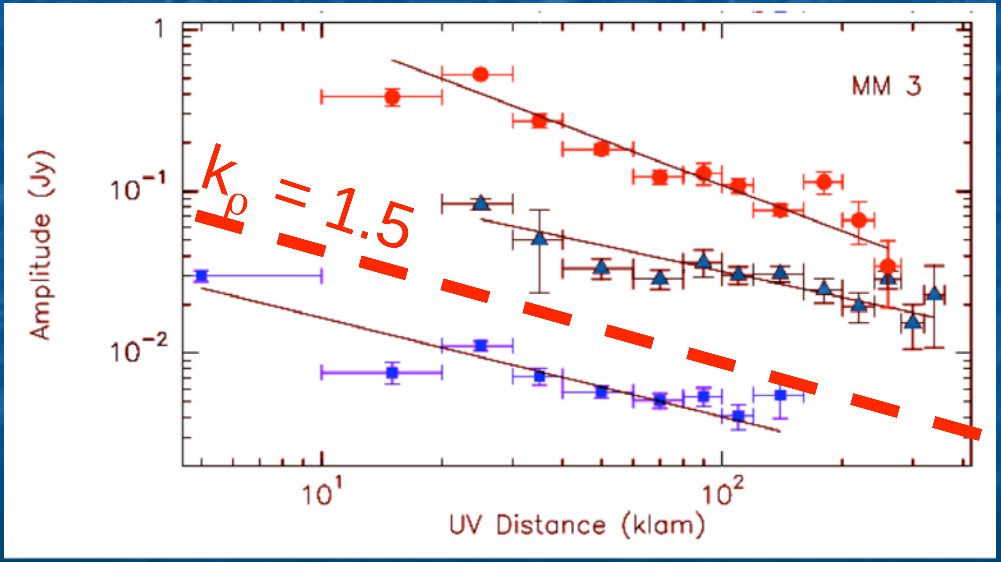


Spitzer/IRAC (left) and Spitzer/MIPS (right), Rathbone et al. (2005)

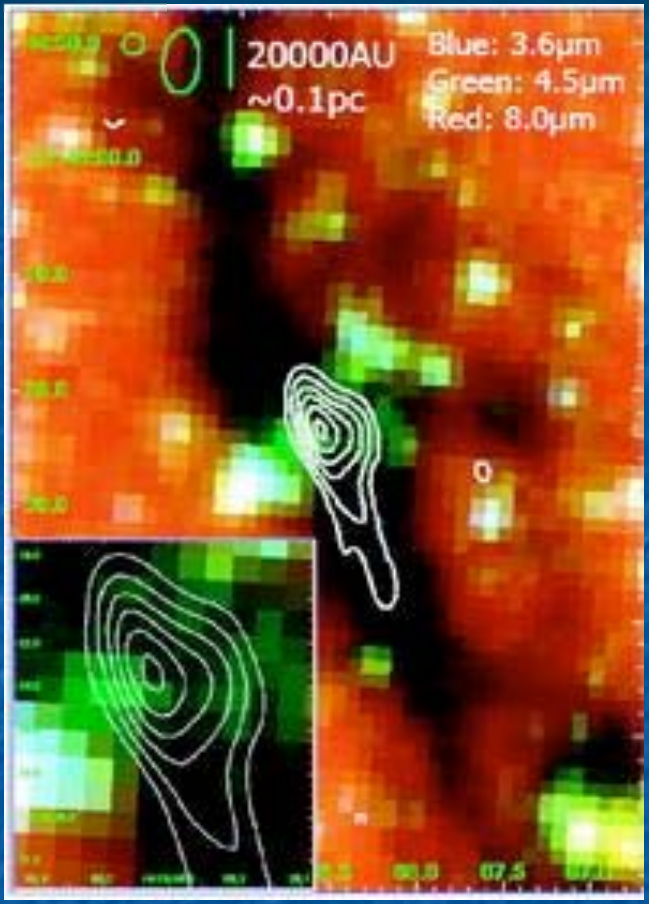
- Massive stars form in gas clumps seen in mm continuum or lines, or in IR absorption (IRDCs)
- Typical properties:
  - $M \sim 10^3 - 10^4\ M_\odot$
  - $R \sim 1$ pc
  - $\Sigma \sim 1$ g cm^{-2}
  - $\sigma \sim$ few km s^{-1}
- Properties very similar to young rich clusters

Massive Cores

- Largest cores in clumps: M ~ 100 $M_\odot$, R ~ 0.1 pc
- Cores have powerlaw density profiles, index $k_\rho \approx 1.5$
- Some are starless



Core density profile in 3 wavelengths, Beuther et al. (2007)

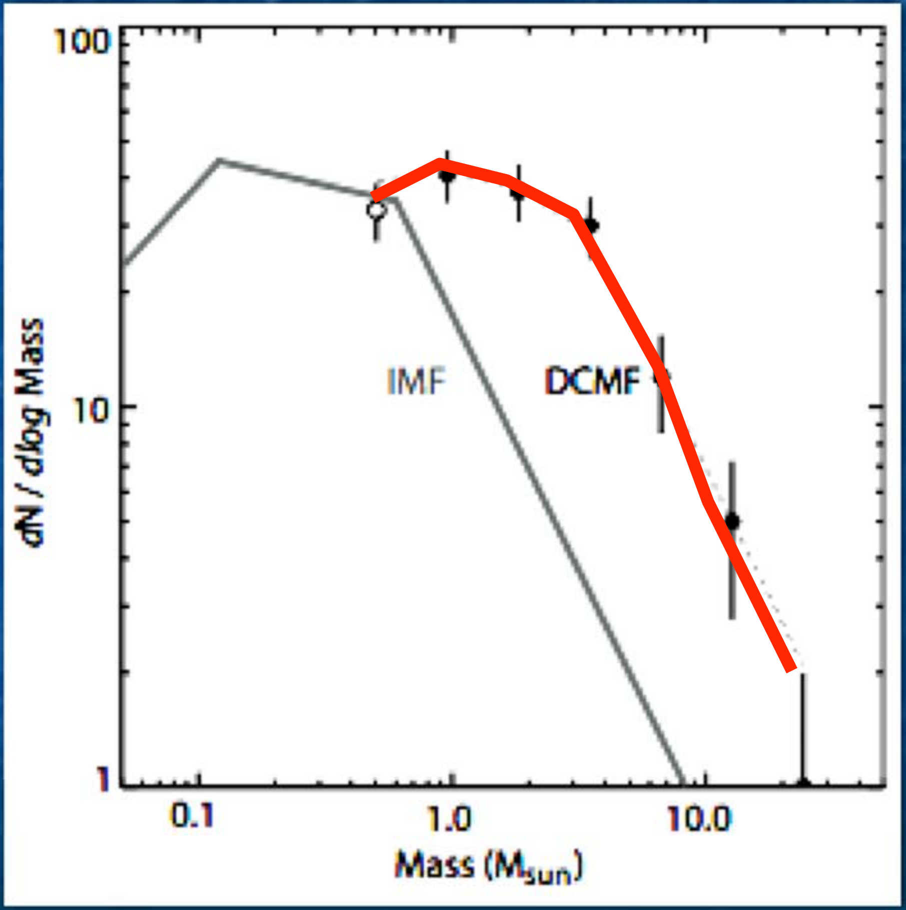


Core in IRDC 18223-3, Spitzer/IRAC (color) and PdBI 93 GHz continuum (contours), Beuther et al. (2005, 2007)

Clue I: The Core Mass Function

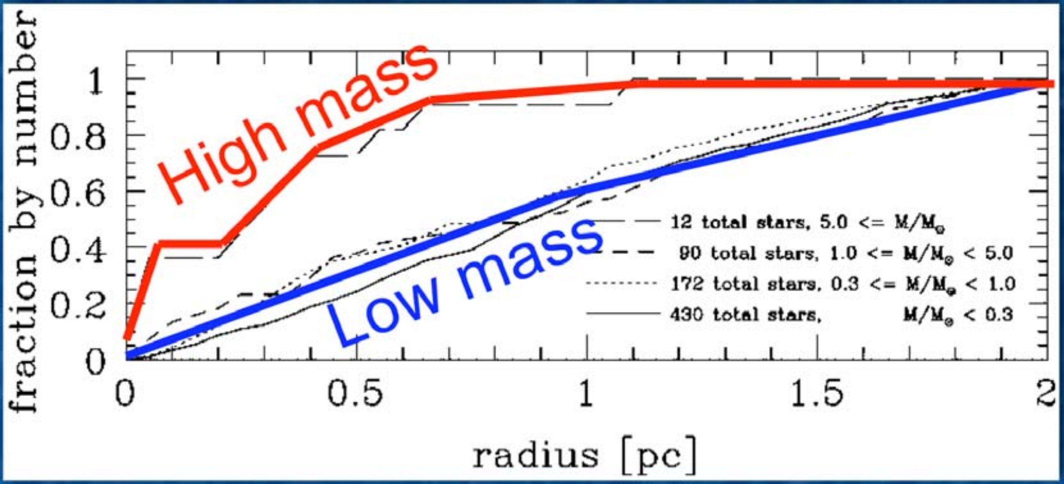
(Motte, Andre, & Neri 1998, Johnstone et al. 2001,
Reid & Wilson 2005, 2006, Lombardi et al. 2006, Alves et al. 2007)

- The core MF is similar to the stellar IMF, but shifted to higher mass a factor of a few
- Correspondence suggests a 1 to 1 mapping from core mass to star mass

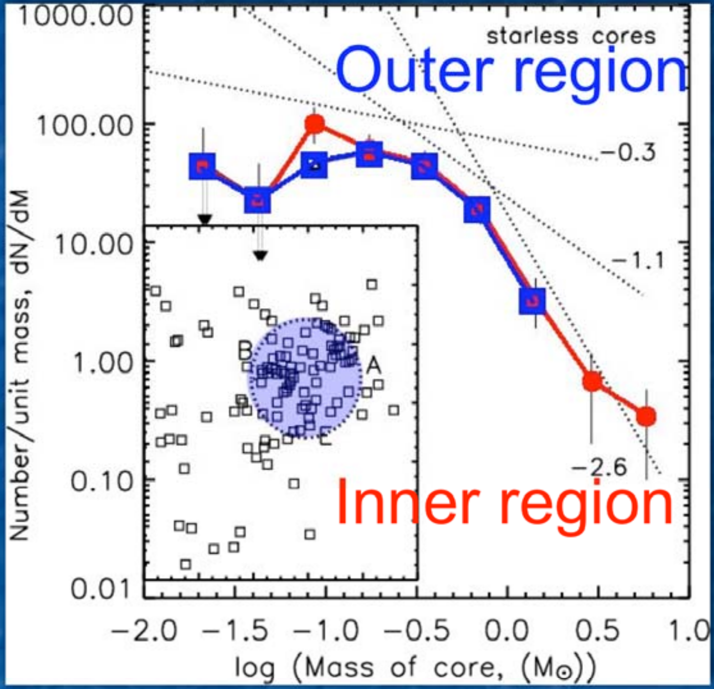


Core mass function in Pipe Nebula (red) vs. stellar IMF (gray) (Alves, Lombardi, & Lada 2007)

Clue II:
Core Spatial Distributions



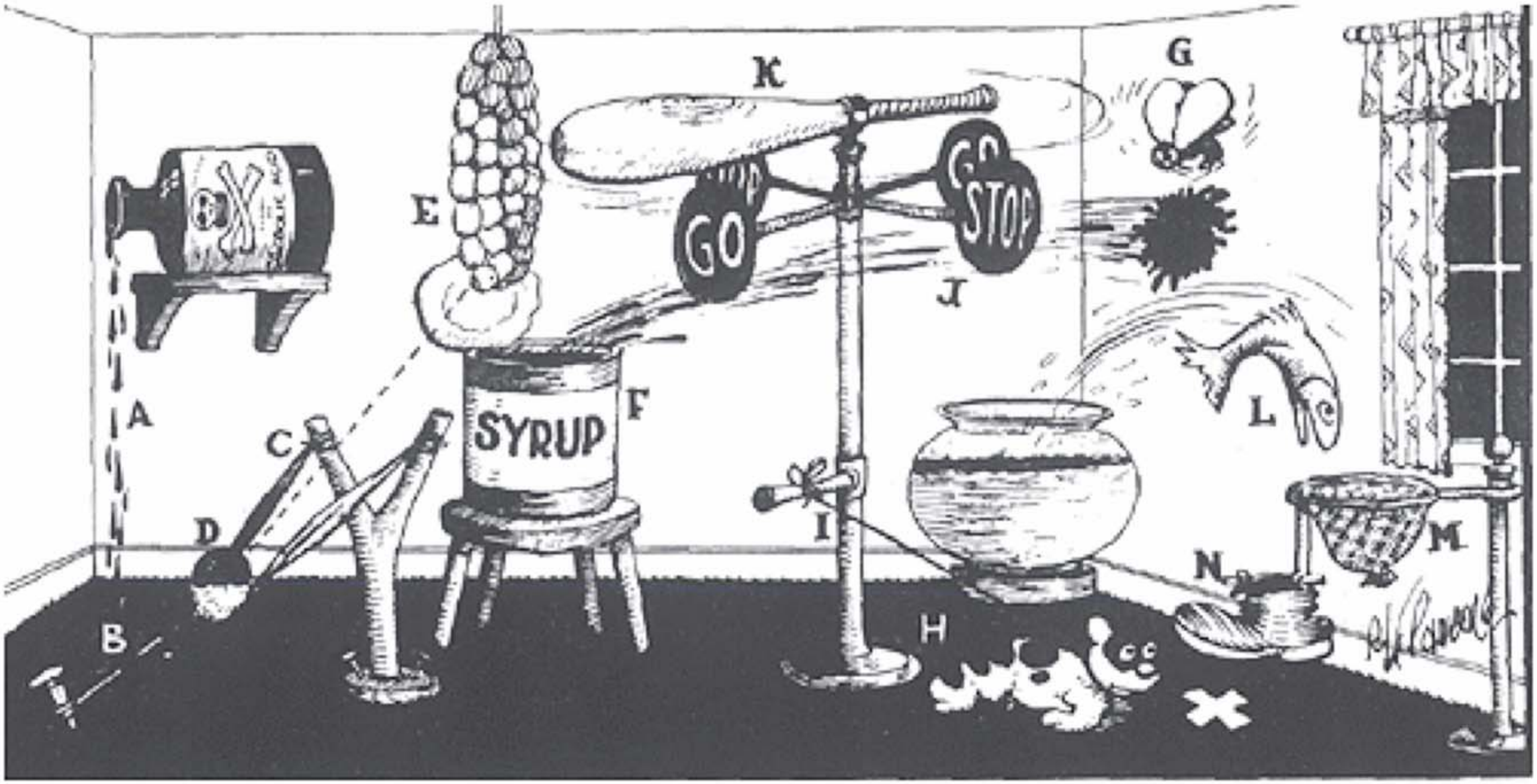
Fraction of stars vs. radius for stars of low mass (blue) and high mass (red) stars in the ONC (Hillenbrand & Hartmann 1998)



Core mass function for inner (red) and outer (blue) parts of ρ Oph, Stanke et al. (2006)

For both stars and cores, the mass function is position-independent at low mass, but high mass objects are only in cluster / clump centers

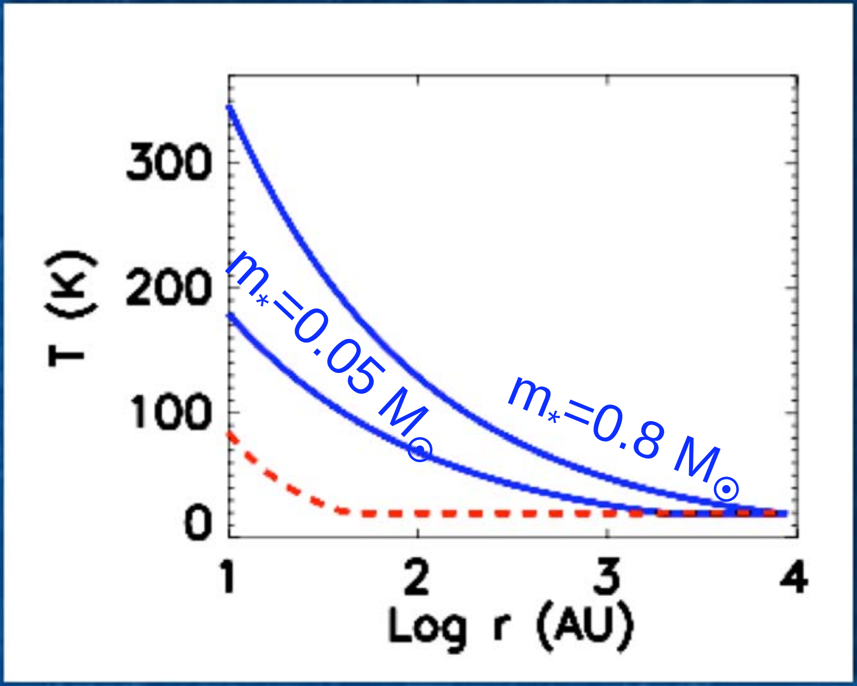
From Core to Star



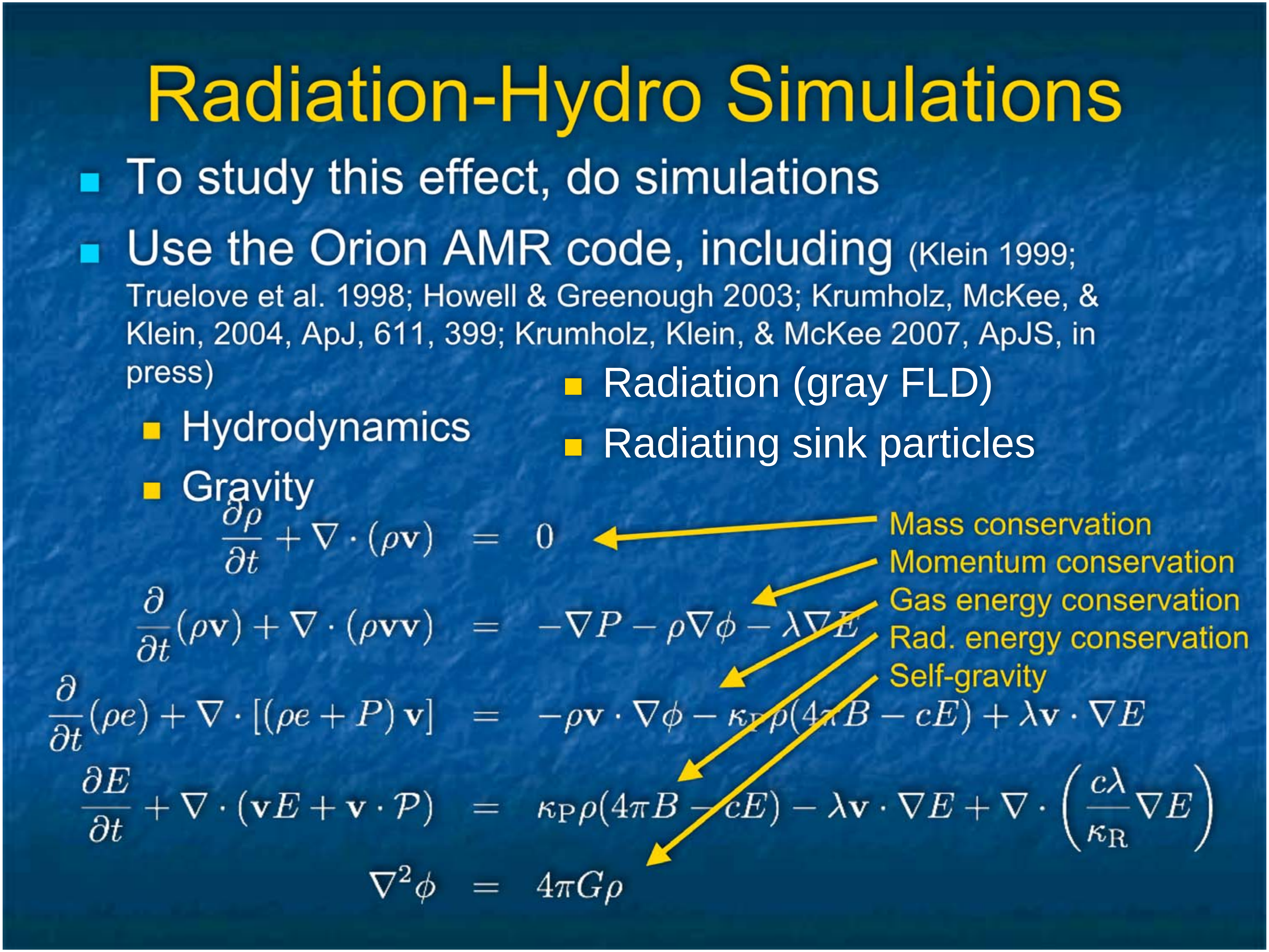
Stage 1: Initial Fragmentation

(Krumholz, 2006, ApJL, 641, 45)

- Massive cores are much larger than M_J (~ $M_\odot$), so one might expect them to fragment while collapsing (e.g. Dobbs et al. 2005)
- However, accretion can produce > 100 $L_\odot$ even when protostars are < 1 $M_\odot$



Temperature vs. radius in a massive core before star formation (red), and once protostar begins accreting (blue)



Radiation-Hydro Simulations

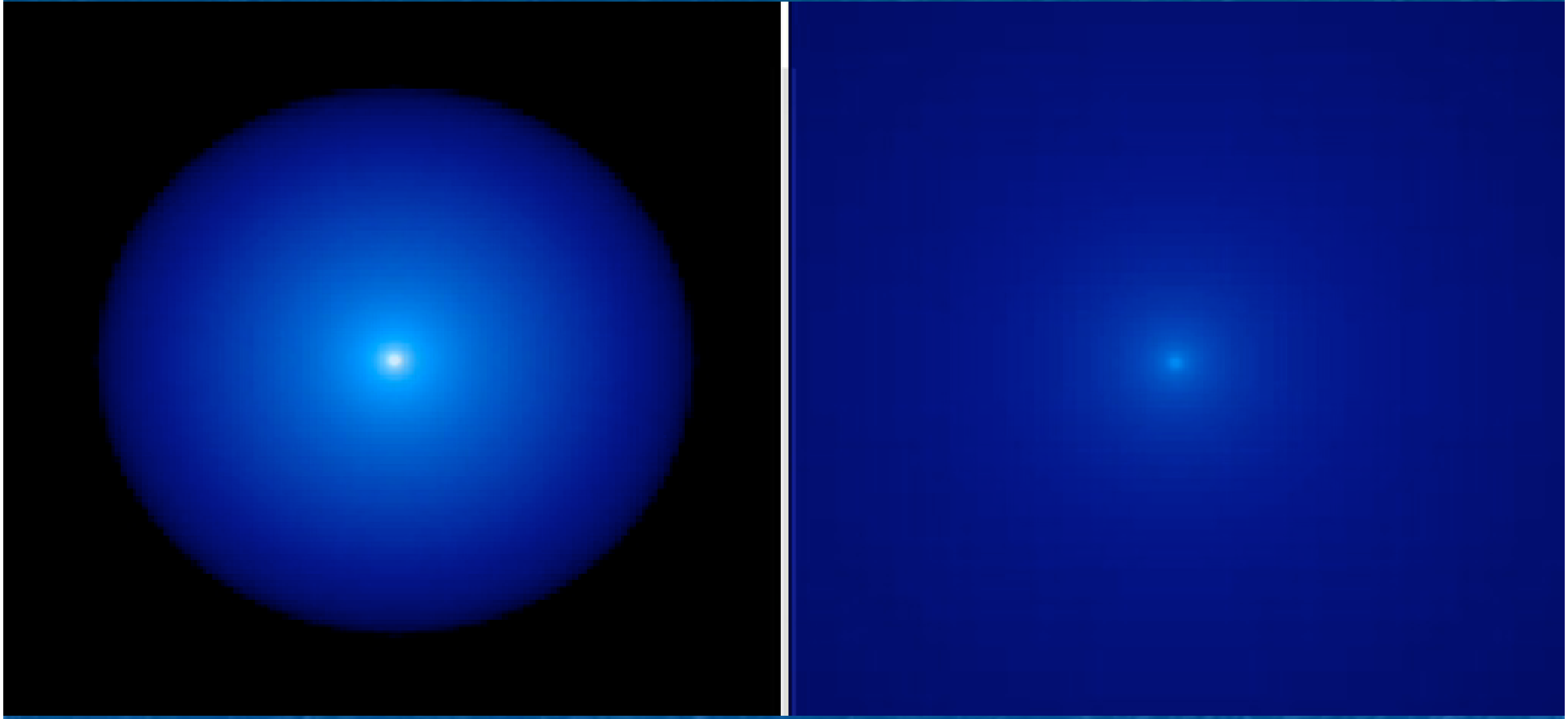
- To study this effect, do simulations
- Use the Orion AMR code, including (Klein 1999; Truelove et al. 1998; Howell & Greenough 2003; Krumholz, McKee, & Klein, 2004, ApJ, 611, 399; Krumholz, Klein, & McKee 2007, ApJS, in press)
  - Hydrodynamics
  - Gravity
  - Radiation (gray FLD)
  - Radiating sink particles

$$
\begin{aligned}
\frac{\partial \rho}{\partial t} + \nabla \cdot (\rho \mathbf{v}) &= 0 \\
\frac{\partial}{\partial t}(\rho \mathbf{v}) + \nabla \cdot (\rho \mathbf{v}\mathbf{v}) &= -\nabla P - \rho \nabla \phi - \lambda \nabla E \\
\frac{\partial}{\partial t}(\rho e) + \nabla \cdot [(\rho e + P)\,\mathbf{v}] &= -\rho \mathbf{v} \cdot \nabla \phi - \kappa_{\mathrm{P}} \rho (4\pi B - cE) + \lambda \mathbf{v} \cdot \nabla E \\
\frac{\partial E}{\partial t} + \nabla \cdot (\mathbf{v} E + \mathbf{v} \cdot \mathcal{P}) &= \kappa_{\mathrm{P}} \rho (4\pi B - cE) - \lambda \mathbf{v} \cdot \nabla E + \nabla \cdot \left(\frac{c\lambda}{\kappa_{\mathrm{R}}} \nabla E \right) \\
\nabla^2 \phi &= 4\pi G \rho
\end{aligned}
$$

Mass conservation
Momentum conservation
Gas energy conservation
Rad. energy conservation
Self-gravity

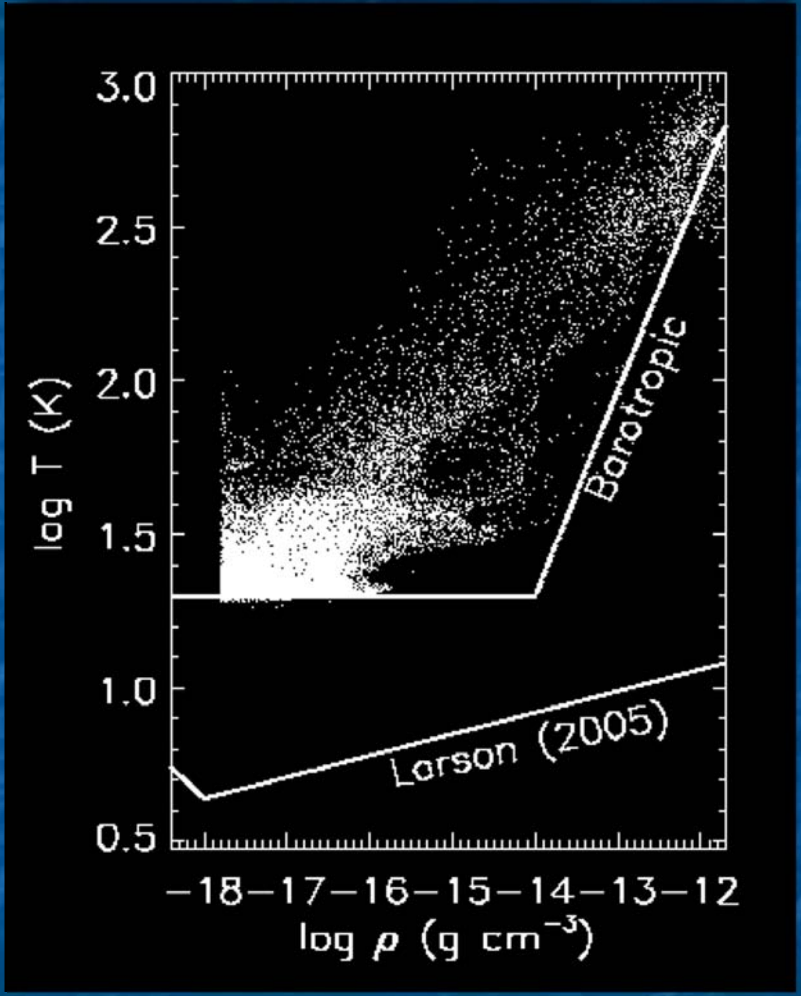
Simulation of a Massive Core

(Krumholz, Klein, & McKee, 2007, ApJ, 656, 959)



- Simulation of 100 $M_\odot$, 0.1 pc turbulent core
- LHS shows Σ in whole core, RHS shows 2000 AU region around most massive star

Massive Cores Fragment Weakly



Column density with (upper) and without (lower) RT, for identical times and initial conditions

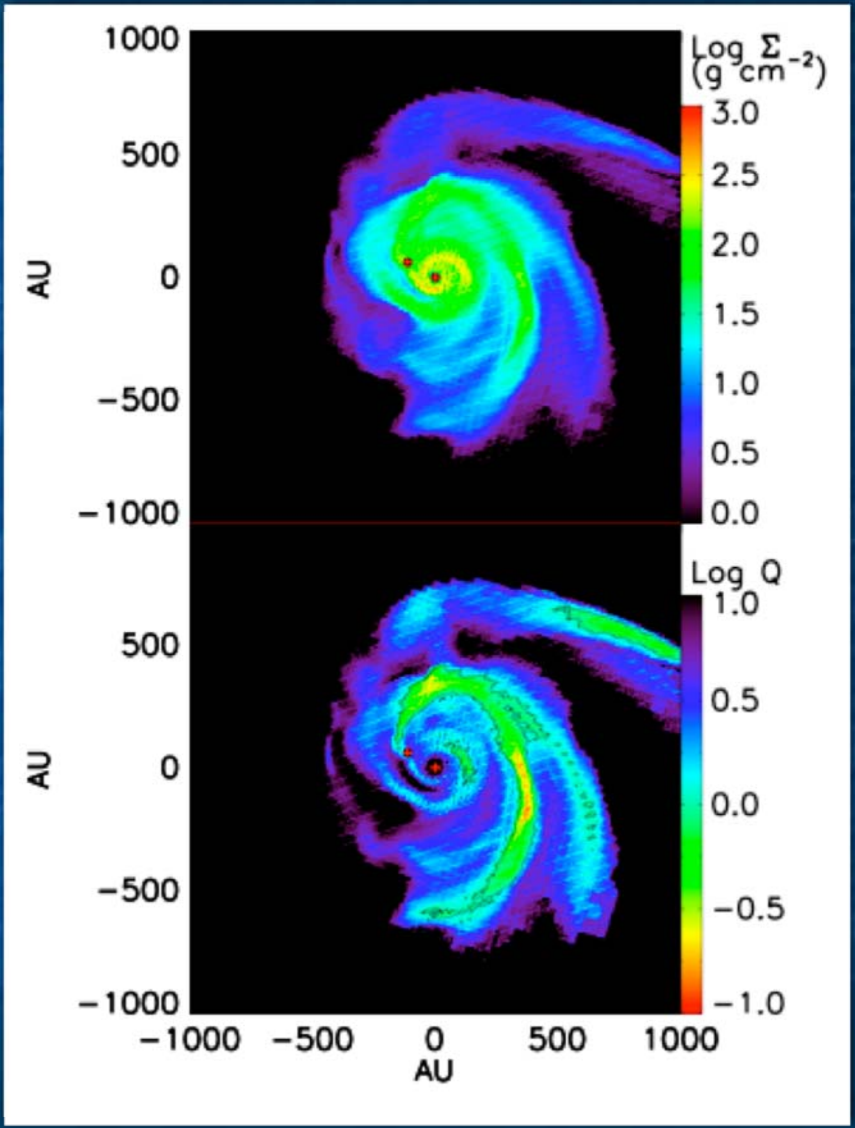
Effective EOS when primary star is < 2 $M_\odot$

- With RT: 6 fragments, most mass accretes onto single largest star through a massive disk
- Without RT: 23 fragments, stars gain mass by collisions, disk less massive
- Conclusion: radiation inhibits fragmentation
- Barotropic or optically-thin cooling EOS fails

Stage 2: Massive Disks

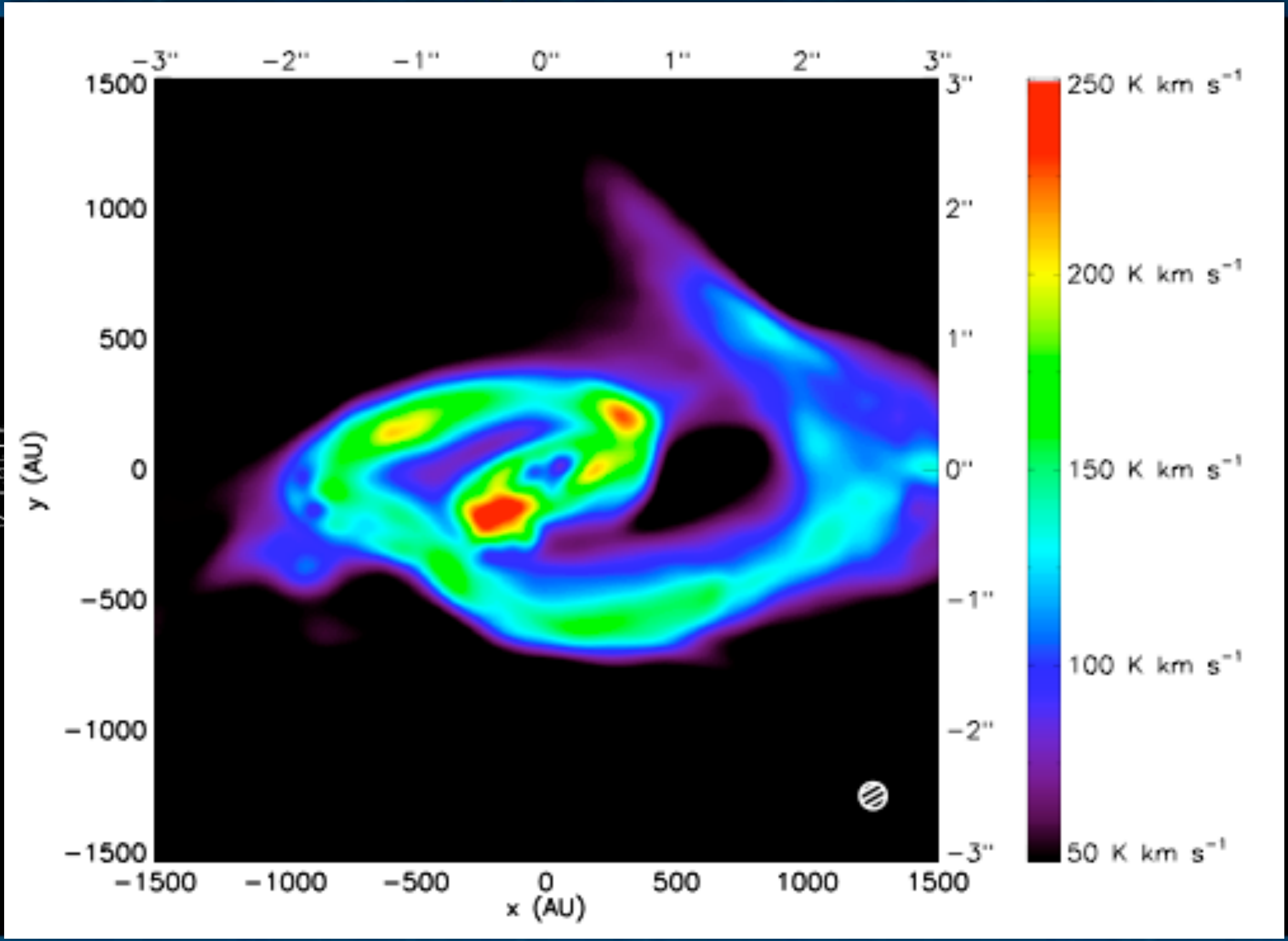
(Kratter & Matzner 2006; Krumholz, Klein, & McKee 2007; Kratter, Matzner & Krumholz, 2007, in prep.)

- $M_{disk} / M_* \approx 0.2 - 0.5$
- Global GI creates strong m = 1 spiral pattern
- Disks accrete very rapidly; $\alpha_{eff} \sim 1$
- Disks reach Q ~ 1, form fragments that migrate inward. Tight binaries likely result.



Surface density (upper) and Toomre Q (lower)

Observing Massive Disks

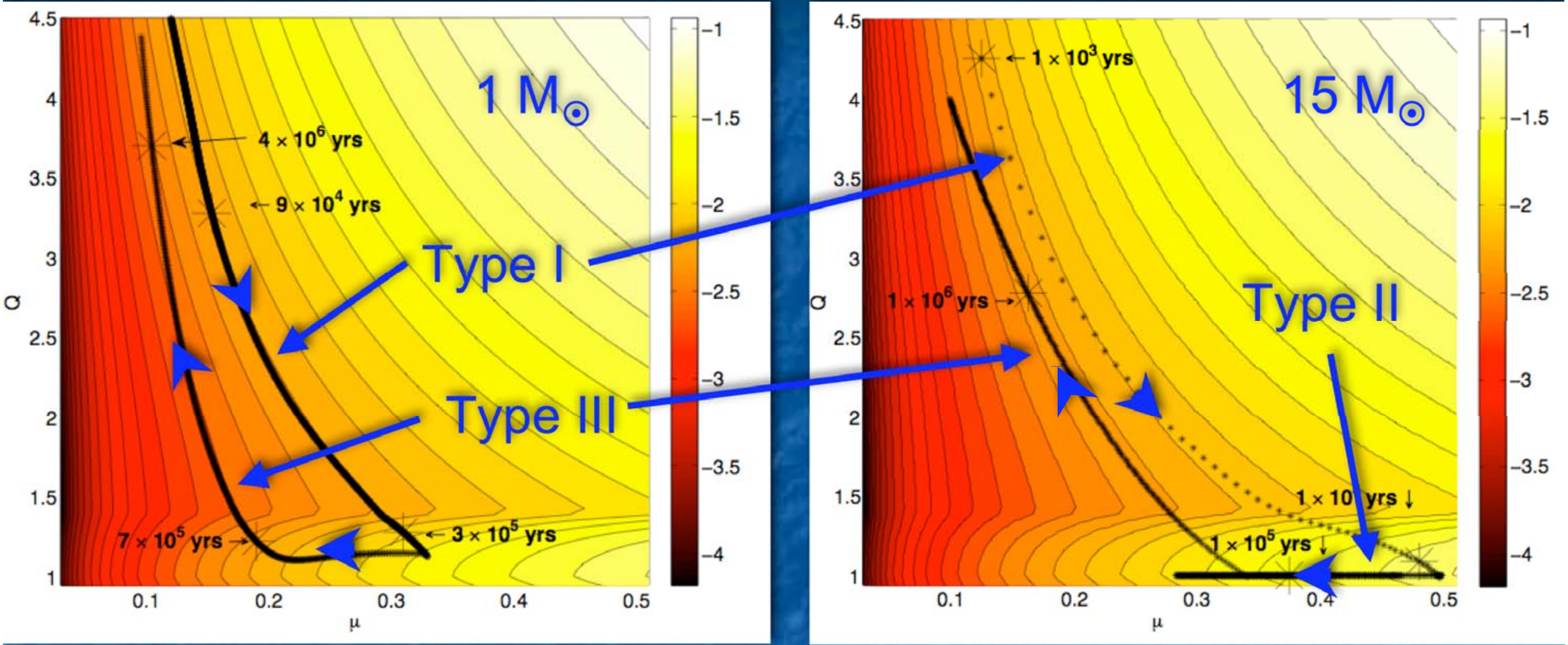


Integrated T_B in simulated 1000 s / pointing ALMA observation of disk at 0.5 kpc in CH_3CN 220.7472 GHz (Krumholz, Klein, & McKee, 2007, ApJ, 665, 478)

Understanding Massive Disks

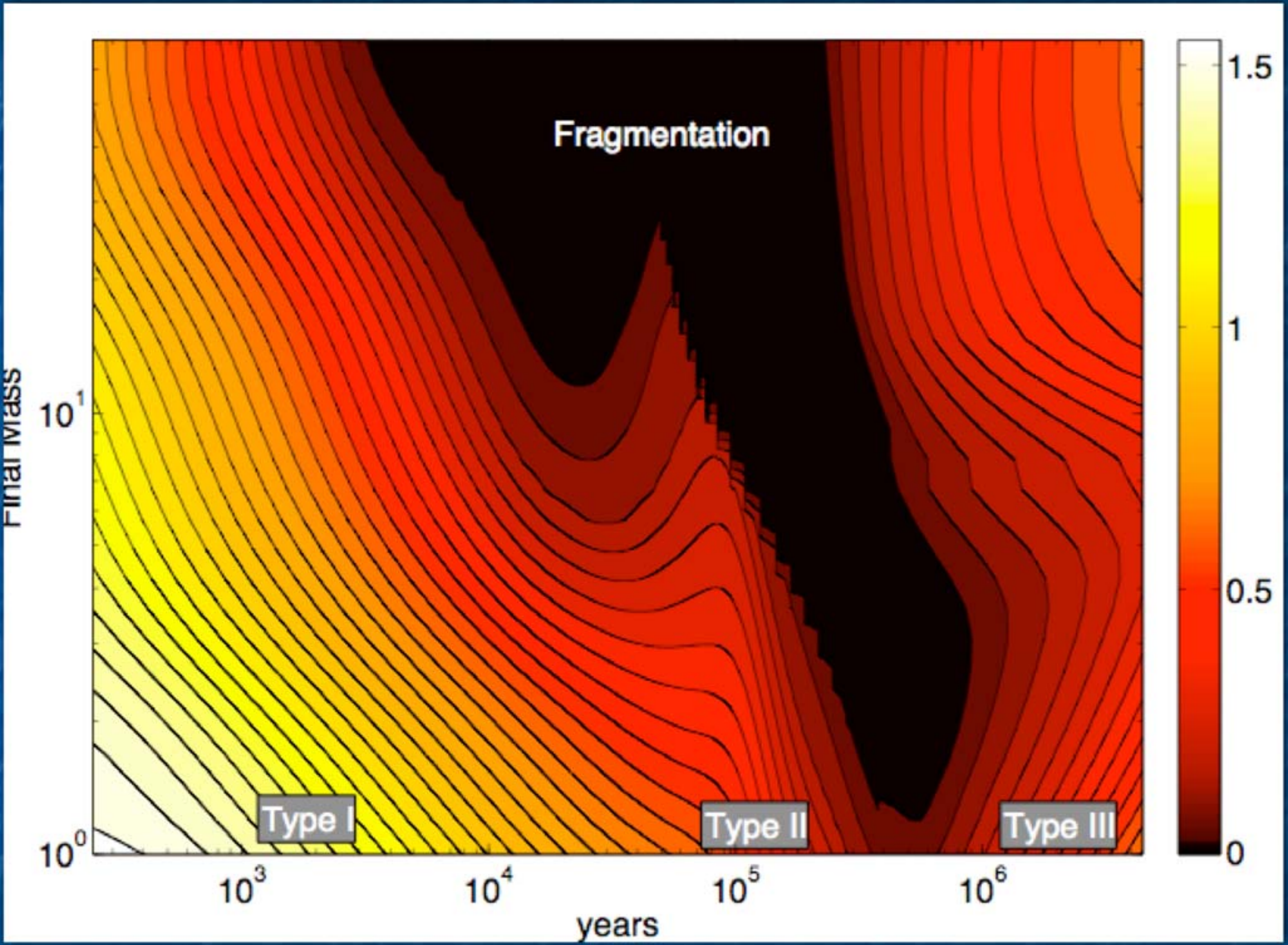
- Accretion rate onto star + disk is $\sim \sigma^3 / G \sim 10^{-3}\ M_\odot / \mathrm{yr}$ in a massive core, but max transfer rate through a stable disk ($\alpha \ll 1$) is $\sim c_s^3 / G \sim 5 \times 10^{-5}\ M_\odot / \mathrm{yr}$ at T = 100 K
- Core accretes faster than stable disk can process $\Rightarrow$ massive, unstable disks
- Study disk evolution using semi-analytic model including accretion, stellar radiation, several ang. mom. transport mechanisms

Model Disk Evolution



Plots show time evolution of disks in the (μ, Q) plane, where $\mu = M_{disk} / (M_{disk} + M_*)$, for 1 and 15 $M_\odot$ stars. The colors and contours show number of orbital periods required to accrete the disk.

Variation in Disk Properties



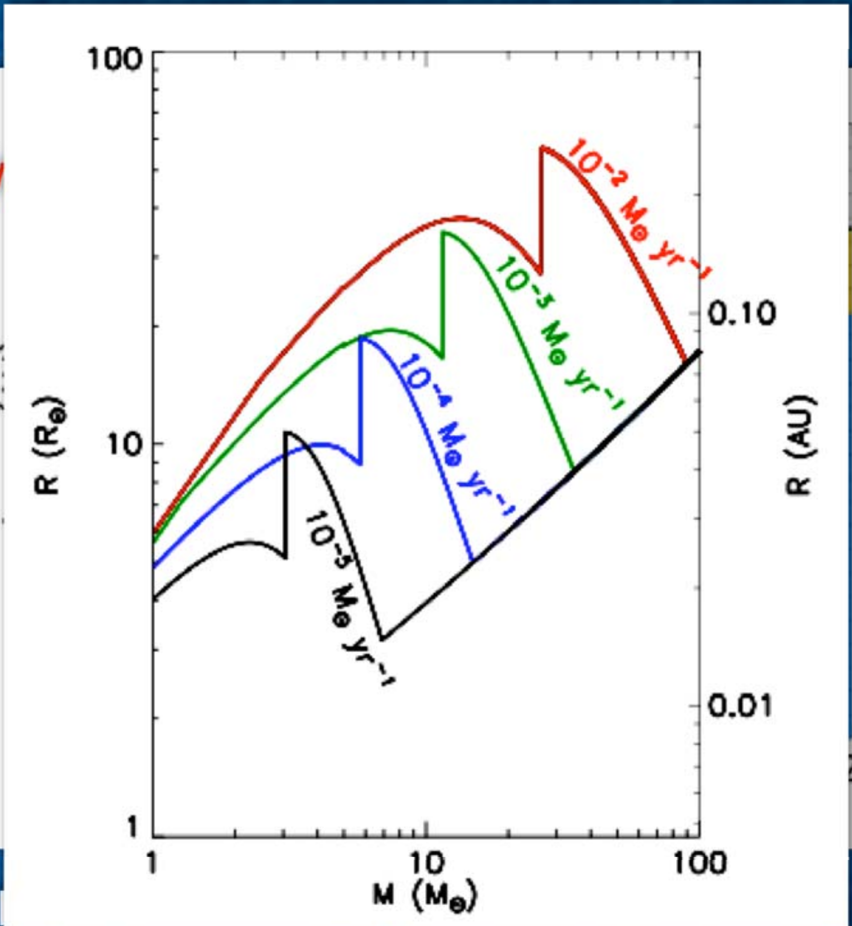
Plot shows Q vs. stellar mass, time.

Implications:

1. Disk fragmentation likely above ~few $M_\odot$ ⇒ explains ubiquity of massive star binaries
2. Large-scale spiral structure present in disks of all stars ~ $M_\odot$ or larger for at least a short period during class 0 phase

Massive “Twins”

(Krumholz & Thompson, 2007, ApJ, 661, 1034)



Radius vs. mass for protostars of varying accretion rates

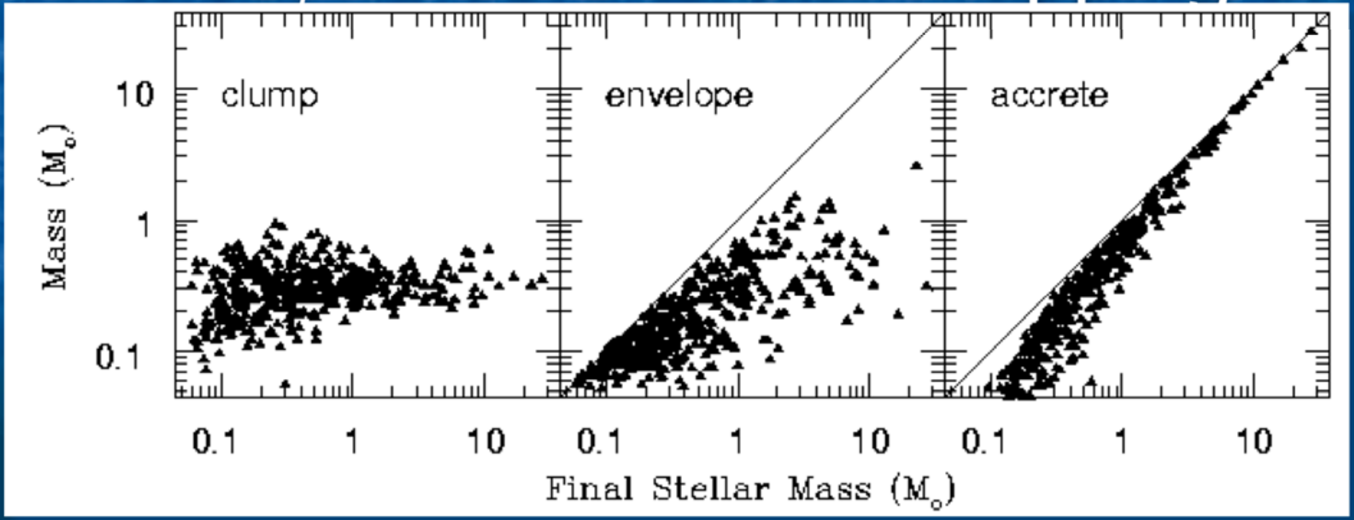
- Massive protostars reach radii ~ 0.1 AU due to D shell burning
- This produces RLOF in close binaries
- Transfer is from more to less massive ⇒ transfer unstable, stabilizes at $q \approx 1$

Stage 3: Radiation Pressure Feedback

See Richard Klein’s talk tomorrow… but the punch line is that radiation can’t stop accretion

Stage 4: Competitive Accretion

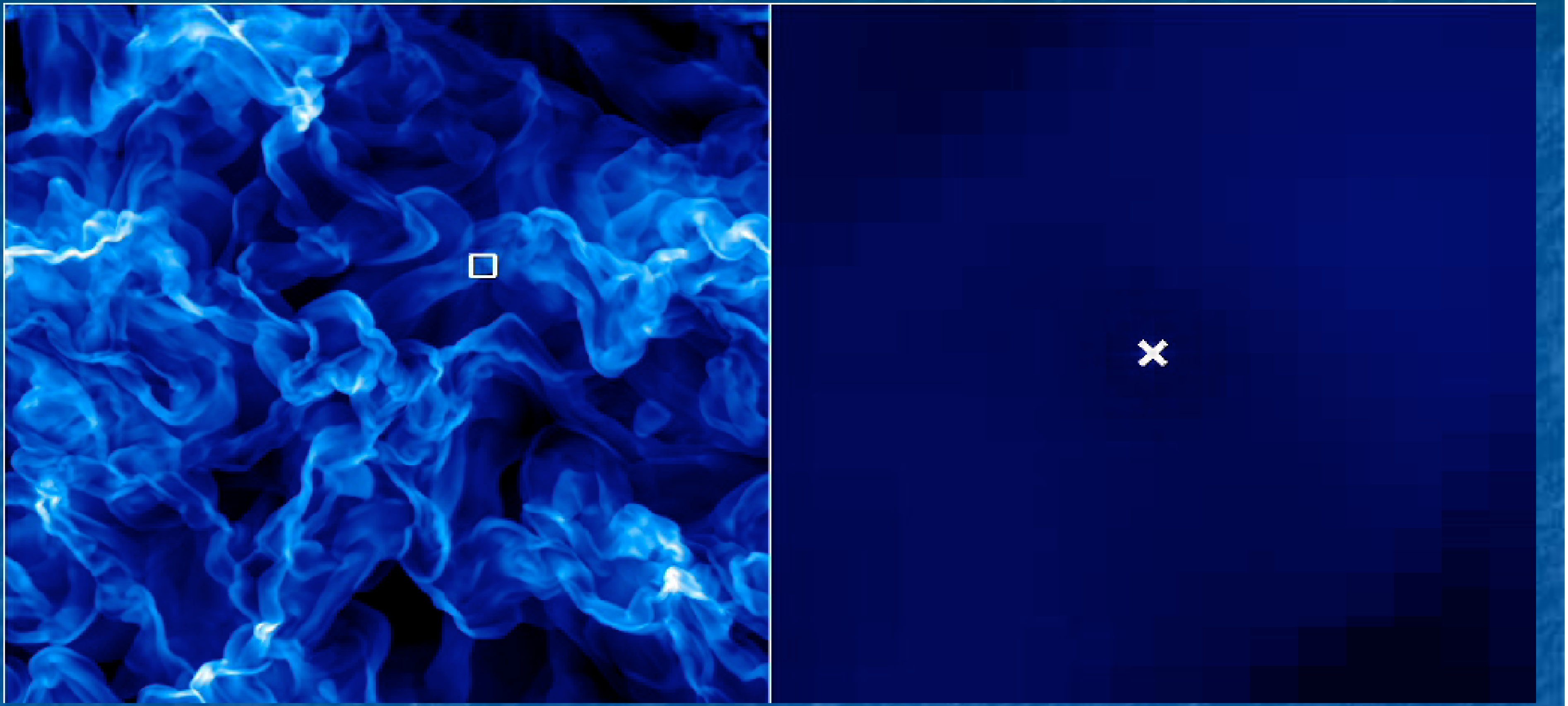
- Once initial core is accreted, could a star gain additional mass from gas that wasn't bound to it originally via BH accretion?
- If so, no core to star mapping exists



Simulation of star cluster formation, Bonnell, Vine, & Bate (2004)

Accretion in a Turbulent Medium

(Krumholz, McKee, & Klein, 2006, ApJ, 638, 369; 2005, Nature, 438, 332)

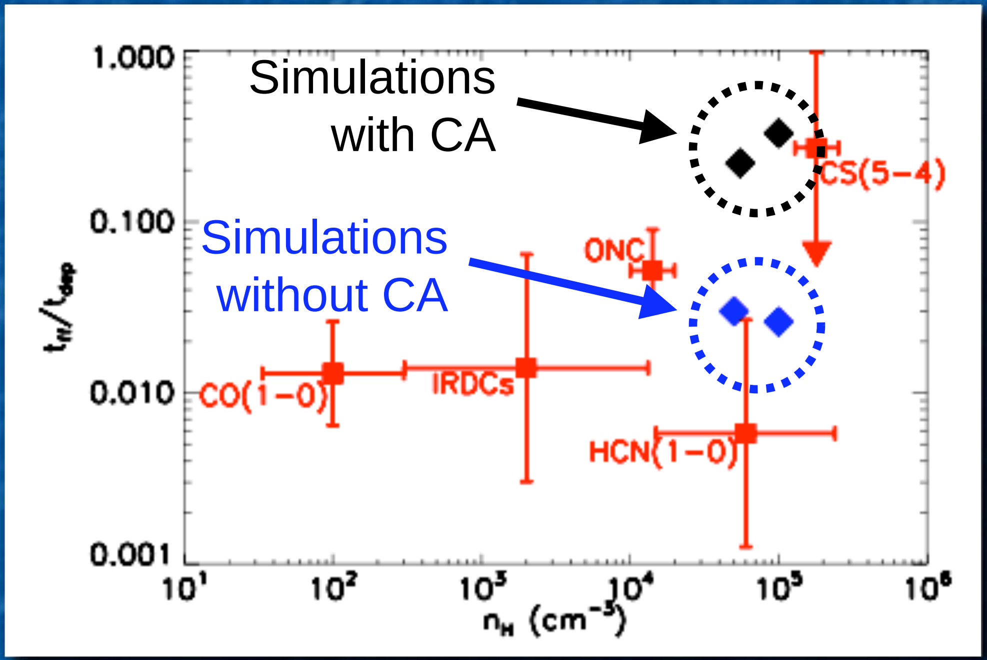


- Result: virialized turbulence ⇒ negligible accretion
- Implication: CA possible only if turbulence decays, cluster collapses to stars in ~1 crossing time

The Star Formation Rate: A Test of Competitive Accretion

(Krumholz & Tan, 2007, ApJ, 654, 304)

- Observe ratio t_{ff} / t_{dep}, in cluster-forming gas clumps (e.g. Gao & Solomon 2004, Wu et al. 2005, Rathborne et al. 2006)
- Compare to ratios from simulations
- CA requires t_{dep} ~ t_{ff}, but observations give t_{dep} ~ 50 t_{ff}



Ratio of free-fall time to depletion time in gas clouds of varying density

Observed SFRs much too low for CA to occur!

Summary

- Massive stars form from massive cores
  - Massive cores fragment only weakly
  - They collapse to massive, unstable disks that form companions
  - Once the core has accreted, the star gains no more mass from elsewhere
- Mass and spatial distributions of massive stars are inherited from massive cores

Finally, thanks to the organizers for giving me a reason to escape New Jersey…